MSE2151
Astronomy Laboratory – Stars

Lab Manual
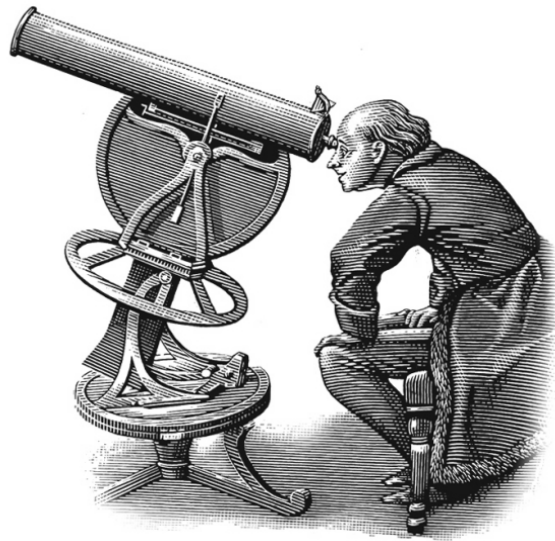
Lab Schedule
MSE2151: Astronomy Laboratory – Stars
Spring 2022

Week 1:   Lab A – Working with Numbers, Graphs & the Computer

Week 2:   Lab B – Energy Transport in Stars

Week 3:   Lab C – Introduction to Spectroscopy

Week 4:   Lab D – Introduction to Optics

Week 5:   Lab E – Building a Galileoscope

Week 6:   Lab F – Classification of Stellar Spectra

Week 7:   Lab G – Effects of the Atmosphere on Astronomical Observations

Week 8:   Lab H – Taking a Star's Temperature

Week 9:   Lab I – Trigonometric Parallax

Week 10:   Lab J – The Distance to a Star Cluster

Week 11:   Lab K – Measuring the Mass of the Andromeda Galaxy

Week 12:   Lab L – Expansion of the Universe

Week 13:   MAKEUP LAB

Observing sessions for the "Observatory Lab" will be scheduled throughout the semester.

Observatory Lab
<u>The Villanova Public Observatory</u>

<u>PURPOSE:</u>

The science of astronomy began with the first humans who looked up into the nighttime sky and wondered "What's going on up there?" This semester, you will explore some of the answers to this question in your lecture course and in this lab course. Most of this exploration will happen in the classroom or in the lab room. This lab will give you the opportunity to go outside and actually look at the nighttime sky, using both the unaided eye and a powerful modern astronomical telescope.

<u>EQUIPMENT:</u>

The Villanova Public Observatory, accessible via the Department of Astrophysics & Planetary Science on the 4th floor of Mendel Science Center.

The Villanova Public Observatory

14" Celestron Schmidt-Cassegrain Telescope

Instructions

To complete this lab you must fulfill 4 requirements:

1. At least once during the semester you must attend an evening observing session at the Villanova Public Observatory, which will be conducted by students from the Department of Astrophysics & Planetary Science. During this session, the students will take you on a tour of the most interesting objects visible in that night's sky.
2. Have one of the Observing Assistants sign and date the cover page of your lab report.
3. Sign that evening's Observatory Attendance Sheet. (If you can't find it, ask one of the Observing Assistants.)
4. Write a 2-page report (double spaced, 11- or 12- point font, normal margins) describing your visit to the Observatory. Include the sky conditions, describe the type of telescope you used, and the objects you viewed. If you have suggestions for improvements of the observing sessions, or any general comments about the experience, please feel free to include them in your report.

The observatory is accessed from the 4th floor of the Mendel Science Center through the Department of Astrophysics & Planetary Science. Once on the 4th floor, the Observing Assistants will direct you to the telescopes.

During the semester, and beginning Tuesday January 18th, the Observatory will be open Monday thru Thursday evenings, initially from 7-9 PM and then, after Daylight Savings Time begins on March 13, from 8-10 PM. The Observatory will not be open during Spring Break or on any University holidays or Snow Days. The last possible night to complete this assignment is the last day of classes Thursday April 28. The Observatory will then be closed for the semester.

Lab reports can be handed in any time during the semester. The last time they will be accepted is 5 PM on the Reading Day, Friday April 29.

There is no Make-up for the Observatory Lab! Since it is one of 13 graded lab assignments this semester, then 1/13 of your grade will be a zero if you don't get to the Observatory! There will be plenty of clear nights this semester, but don't wait til the end. It is entirely possible that the last week or so could be clouded out or rained out – it's happened before.

Frequently Asked Questions

What will I be looking at?

It depends on what is available in the sky. The Moon looks great through our telescopes and the best time to view it is a few days after the First Quarter phase. In this phase, the Moon is high in the sky around the time of sunset and the side illumination creates very pronounced shadows on the lunar surface, allowing the rugged terrain to be seen clearly. During the Spring 2022 semester, First Quarter phases occur on **Feb 8**, **Mar 10**, **Apr 9**, and **May 9**.

This will not be a great semester for planet viewing. Mars, Jupiter, and Saturn will all be in the daytime sky throughout the semester and thus not viewable for the Observatory Lab. Venus will be the "morning star" this semester, visible by eye above the eastern horizon shortly before sunrise. Saturn, Mars, and Mercury will be in the same area of the sky as Venus and so – if you want to get up early − you may be able to catch sight of three or four planets before you go to breakfast. If you'd like to keep track of the visibility of the planets throughout the year, check out: https://www.timeanddate.com/astronomy/night/

In addition to the Moon, there will be plenty of other objects available to keep the telescope busy, including star clusters, bright nebulae, and the Andromeda galaxy.

If you'd like to keep informed about upcoming celestial events in general, check out https://www.timeanddate.com/astronomy/sights-to-see.html

Is the Observatory Open tonight?

If "tonight" is a Monday thru Thursday and a regular class day, then the chances are good. It all depends on the weather. What is the best way to tell? Look up shortly after sunset. If you can see the stars, we're probably open. If you want a more high-tech solution, you can go to one of the many weather websites (or apps) available. Most of these show cloud cover and some will project the cloud cover a day or so in advance. One such site is http://www.weatherstreet.com/weather-forecast/Villanova-PA-19085.htm.

What if I don't fulfill all 4 of the requirements listed in the Instructions?

No credit will be issued for the Observatory lab and a grade of "0" will be recorded.

Astronomy Laboratory - Planets
Mendel Science Experience 2151
Spring 2022

Observatory Lab
<u>The Villanova Public Observatory</u>

Name: _____

Lab Section: _____

Date: _____

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Observatory Assistant

Name: _____

Signature: _____

Date: _____

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Lab A
Working with Numbers, Graphs, and the Computer
Adapted from: *Laboratory Experiments in Astronomy*, Johnson & Canterna

PURPOSE:

To become acquainted with the Excel spreadsheet and to review some mathematical concepts that will be implemented in upcoming labs.

EQUIPMENT:

The computer and the Excel spreadsheet.

Introduction

By their nature, the sciences are quantitative. Science involves making measurements, analyzing quantitative relationships, and making predictions based on mathematical models. In this exercise we will review some of those mathematical fundamentals that will be used in subsequent exercises. This review will help you pinpoint any areas you might need to review in more depth. The magnitude scale, a strictly astronomical concept, will be introduced. This has been the shorthand notation used for centuries to describe the brightness of stars.

Lab Procedure

PART 1: SCIENTIFIC NOTATION

Astronomy is a science which deals with the very large and the very small. For example, the distance to the nearest star (beyond the Sun) is approximately 39,900,000,000,000 km, while the distance between the electron and proton in an atom of Hydrogen in interstellar space is about 0.00000000529 cm. It can be cumbersome to deal with such extreme numbers and a more compact way of dealing with them is to express them in **scientific notation**.

The number 13,700,000 is expressed in scientific notation as 1.37×10^7. "10^7" is shorthand for 1 followed by 7 zeros, $10 \times 10 \times 10 \times 10 \times 10 \times 10 \times 10$ (10 times itself 7 times), or 10,000,000. 1.37×10^7 equals 13,700,000 or 1.37×10 million. Any number can be expressed as a number between 1 and 10, times 10^n. Take 207,000 as an example. There are 5 decimal places to the right of the 2. Thus $207,000 = 2.07 \times 10^5$. Multiplication by 10 is equivalent to shifting the decimal point in a number one place to the right. Multiplication by 10^n is equivalent to shifting the decimal point n decimal places to the right, or to multiplying by 10 n times.

This technique is also useful for numbers smaller than 1. For example, 0.0012 can be represented as 1.2×10^{-3} (or 1.2 divided by 1000). A negative exponent indicates how many places the decimal is shifted to the left. 10^{-3} is equivalent to 0.001 or $1/10^3$. *Note:* A negative exponent does <u>not</u> mean that the number itself is negative; only that it is between 0 and 1.

EXERCISE 1-1: *Using Table 1 included on the last page of this lab, convert the numbers given in the first column ("Standard Notation") into scientific notation.*

EXERCISE 1-2: *Create a single-column Excel spreadsheet table using the "Standard Notation" numbers from Table 1. Label the top row of the column (you can call it "Standard Notation.") Your numbers should look <u>exactly</u> as they do in Table 1. You will need to adjust the number of decimal places displayed for each number. To do this, first highlight a cell, then right-click and select **Format Cells**. Select **Number** and type in the desired number of decimal places needed to match the format in the Table 1.*

EXERCISE 1-3: *On your Excel spreadsheet, copy your column of numbers and paste it into the adjacent column (i.e., column B). Now have Excel convert these numbers to scientific notation for you. To do this, highlight the cells, right-click, and select **Format Cells**. Then choose **Scientific Notation** and type in the desired number of decimal places (in this case, 2). Label this column "Scientific Notation."*

Notice that, in Excel, the format used for scientific notation is somewhat different from that described above. E.g., the quantity 2.07×10^{-5} is written by Excel as 2.07e-5, where the "e" stands for *"times 10 to the"*. <u>This notation must be used to enter scientific notation into Excel.</u>

Also notice that it is easier to keep track of **significant figures** in scientific notation. Simply, all the figures used in scientific notation are significant. Thus, the number 2.07e-5 is implied to have three significant figures and the number 2.070e-5 is implied to have 4 significant figures. In this last case, the trailing zero is significant and thus a result expressed as 2.070e-5 can be assumed to be more precise than a result expressed as 2.07e-5.

PART 2: LOGARITHMS

Another shorthand way of expressing very large or very small numbers is with **logarithms**. The **logarithm** or **"log"** of a number is simply the exponent to which 10 must be raised to equal that number:

$$x = 10^y \quad \Leftrightarrow \quad y = \log(x) \tag{1}$$

For example, the number one hundred million (100,000,000) can be written as 10^8 and, therefore, **log**(100,000,000) = 8.

Negative numbers do not have logarithms since there is no power to which 10 can be raised that will yield a negative number. However, the logarithm itself can be negative, and corresponds to a number greater than zero but less than one. The log of a number can also be a non-integer. Consider **log**(250). The number 250 is greater than 10^2 (i.e., 100) but less than 10^3 (i.e., 1000) and so we would expect its logarithm to be between 2 and 3. In fact, **log**(250) = 2.40.

EXERCISE 2-1: *In column C of your Excel spreadsheet, compute the logarithms of the numbers in the first column. For example, to compute the log of a number located in cell A2 of your spreadsheet (i.e., column A, row 2) and place the result in cell C2, simply type "=log(A2)" in cell C2. The log of the number in cell A2 will then appear in cell C2. Do this for all entries in column A. Be sure to label the column and adjust the number of decimal places in the displayed logarithms to preserve the number of significant figures in the original numbers.*

NOTE: Logarithms are never expressed in scientific notation. Always use standard notation.

NOTE: With logarithms, only the numbers to the right of the decimal point are significant digits. For example, the number 4,567,000,000 has 4 significant digits. It should be expressed in scientific notation as 4.567×10^9 (again, with 4 significant digits) and logarithmically as 9.6596. In his latter case, the integral portion of the logarithm, "9" actually refers to the power of 10 in the scientific notation form. The decimal portion ".6596" carries the information contained the significant digits.

PART 3: LINEAR RELATIONSHIPS

In this exercise we will see how physical relationships may be examined graphically. In the figure below, we have plotted the weights (on the "y" axis) versus the heights (on the "x" axis) for the Super Bowl-winning 1985 Chicago Bears football team, using the small diamond symbols. It is clear from the plot that there is some general correlation between the heights and weights, i.e., "the taller, the heavier." We have drawn a straight line through the data to indicate a more specific *linear relationship* between the height and weight in this group. The line is intended to represent the **general trend** in the data, while ignoring specific deviations of individuals. In this case, the line appears to give a good fit to the general trend of the data and it is reasonable to conclude that height and weight are linearly-related for this dataset.

Note the outlying point far above the main relationship; these data correspond to William "the Fridge" Perry, one of the first 300 pounders in football and who clearly did not fit into the general height-weight relationship of his teammates!
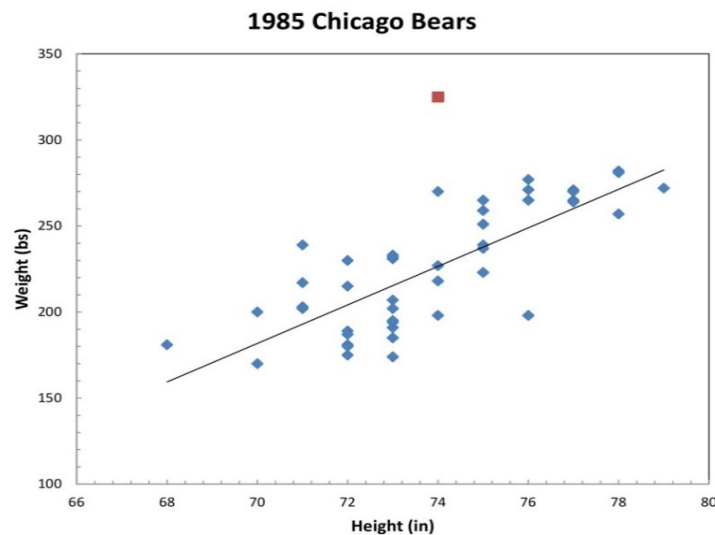
In the Table below we have listed the heights and weights for a different team, namely, a selection of 15 typical college students.

Typical College Students			
Height (inches)	Weight (pounds)	Height (inches)	Weight (pounds)
60	95	69	150
70	173	64	115
67	125	62	112
75	180	64	127
61	105	68	155
73	200	72	160
67	140	65	127
71	150	…	…

EXERCISE 3-1: *Transfer these data to a new Excel spreadsheet. To open a new sheet, click **Sheet2** at the bottom of your current spreadsheet. (If there is no **Sheet2**, click the circled "+" sign to create one.) In your spreadsheet, use only two columns – height and weight. Label the columns appropriately. Be sure that the number format in your spreadsheet matches that of the table.*

EXERCISE 3-2: *Now create a graph of the college students' Weight vs. Height on your spreadsheet. Follow the instructions given in the "**Microsoft Excel Tutorial**" in Appendix I of this Lab Manual. Be sure that the data in the table are labeled and that the plot is properly scaled and annotated.*

Once again, you should observe a nearly linear relationship between the height and weight of people in the group, as you did for the football players in Problem 1. That is, the heights and weights should fall close to a straight line drawn through them.

EXERCISE 3-3: *Using the **trendline** feature of Excel, show the linear relationship that best fits the data. Make sure that the formula for the trendline is displayed clearly on the plot. (Right-click somewhere on the trendline, then click "Format Trendline…", then check "Display equation on chart.").*

Question 3-1: *Does a visual inspection of your plot confirm that the heights and weights for typical college students are linearly related? Discuss the possible causes of the observed relationship. Why might individuals deviate from this simple relationship?*

Question 3-2: *One of the most powerful uses of a graph is that of prediction. Say, for example, that you know the height of a college student but not his/her weight. We can make an "educated guess" and predict the weight of the student by using our graph and the trendline which represents the general relationship between height and weight. Suppose that this person is 6' 2-½" (i.e., 74.5") tall. What weight is predicted for a 74.5" tall college student? Use the formula for your trendline to make this determination. Show all the work in your calculation. Remember your significant digits!*

Question 3-3: *What differences do you find between the heights and weights of the groups shown in Figure 1 (the Chicago Bears) and on your graph (for college students)? What might be the causes of these differences? Do the lines drawn through the data in both graphs look like they are pieces of the same overall relationship? Explain.*

PART 4: NON-LINEAR RELATIONSHIPS

Not all physically interesting relationships are linear. For example, the table below shows the annual amount of U.S. Government expenditures from 1850 to 2010 versus year which, you will shortly see, reflects a very non-linear relationship.

U.S. Government Expenditures			
Year	**$**	**Year**	**$**
1850	29,000,000	**1960**	97,284,000,000
1860	78,000,000	**1980**	590,941,000,000
1880	304,000,000	**1990**	1,252,990,000,000
1900	629,000,000	**2000**	1,788,950,000,000
1920	6,785,000,000	**2010**	3,457,080,000,000
1940	10,061,000,000	…	…

EXERCISE 4-1: *Open a new Sheet in Excel by clicking the **Sheet3** tab and enter the data from the table above. (In your spreadsheet, use only two columns – year and expenditures.) Label the columns appropriately. If the width of the columns in your spreadsheet is too small, there might not be enough room for Excel to write in all the digits in the numbers and you will see "#########" instead. If you see this, simply drag the right edge of the column letter that represents the expenditure data to expand the width of these cells. Be sure to format the cells so that the numbers appear exactly as in the Table above.*

EXERCISE 4-2: *Now create a plot of expenditure (y-axis) versus year (x-axis). Don't forget to scale and label the plot axes appropriately. Find a trendline which seems to represent the data. Note that a linear trendline will definitely <u>not</u> work! Excel offers you several other options (e.g., "polynomial", "exponential", "power law", etc. Find the one which best represents the data and show it on your plot. Make sure to include the formula for the trendline on the plot. What kind of trendline did you use?*

(Hint: Maybe this gives meaning to the phrase:
"Our national debt is increasing at an exponential rate!")

EXERCISE 4-3: *In your Government Expenditures Excel spreadsheet, copy the contents of column B (Expenditures) to column C. Label this column "$ in Scientific Notation." Now use the* **Format Cells** *command to convert the expenditures in column C to scientific notation. Make sure to specify the correct number of significant digits by adjusting the numbers of decimal places displayed.*

EXERCISE 4-4: *In your Government Expenditures Excel spreadsheet, create a "Log $" column in column D to display the expenditures in logarithmic form. Again, be sure the correct number of significant digits is displayed.*

Your plot of Government Expenditures is interesting but notice how hard it is to distinguish among the values from earlier than about 1960. Also, you should see that it would be very difficult to use the graph to predict future spending with any confidence. Both problems might be reduced by looking at the logarithm of the expenses…

EXERCISE 4-5: *Now create a plot of the logarithm of the expenditures in dollars as a function of year. Adjust the scales to fill the plotting space (if necessary) and add labels and a title. Notice that the very non-linear relationship between Year and Expenditure seen in Exercise 4-2 has become much simpler looking. Because of this, we would say that expenditures and time have a* **logarithmic relationship**. *Create a linear trendline in your Year vs. Log($) plot. Does it fit the observed relationship? Don't forget to put the formula for the trendline on the plot.*

Question 4-1: *What does your plot from Exercise 4-5 suggest that the annual expenditure will be in the year 2022? Use the formula for the trendline you just created in Exercise 4-5 to make this prediction. Show all work in your calculation. Note that the formula will return the logarithm of the predicted expense. Record this and convert it to the expense itself using Equation 1. Express the result in both standard notation and scientific notation. Can we be confident in this prediction? Explain why or why not.*

PART 5: MAGNITUDES - AN ASTRONOMICAL EXAMPLE

Astronomers use a logarithmic scale, called the **Apparent Magnitude Scale**, to quantify the brightness of stars as seen from the Earth. These apparent brightnesses depend on both the intrinsic energy output of the stars and on their distances from the Earth. Over 2,000 years ago the Greek astronomer Hipparchus catalogued the positions and brightness values of approximately 6,000 stars visible to the naked eye. The apparent brightness was described by a unit called a **magnitude**, with first magnitude being the brightest visible stars and sixth magnitude corresponding to the faintest. Contrary to our intuition, the bigger (i.e., more positive) the apparent magnitude, the fainter the object. The brightest object in our sky – the Sun – has the most negative apparent magnitude. See the figure below.
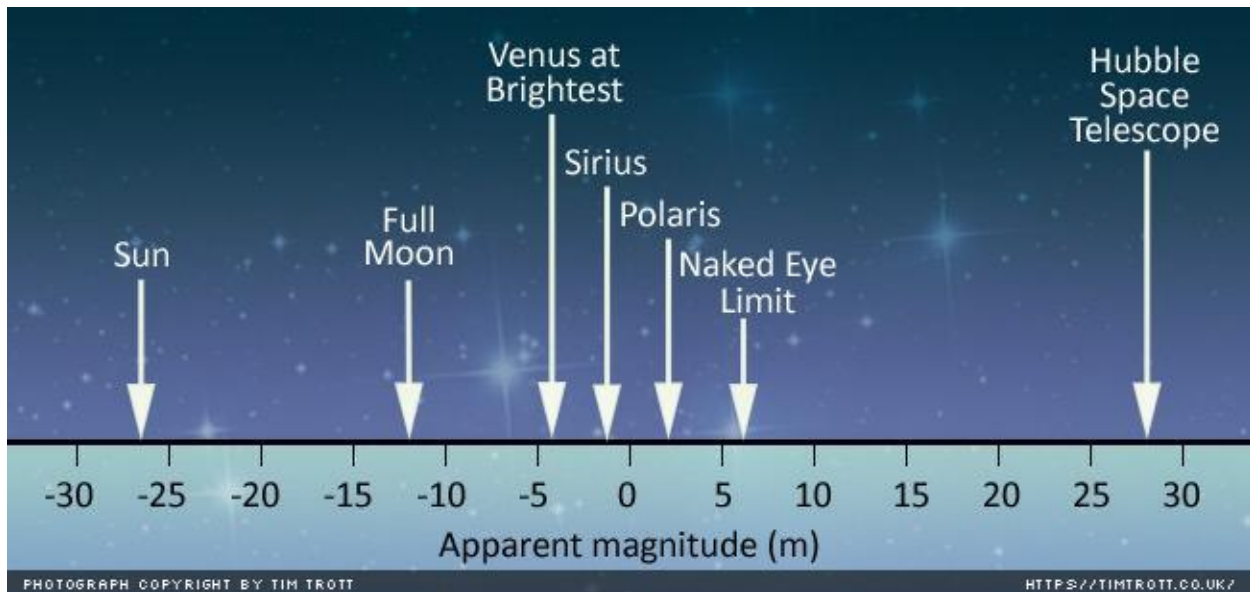
In the following Table we have listed the apparent brightness relative to the star Mizar for the ten stars that are closest to the Sun. For example, the star Proxima Cen appears 2.54×10^{-4} times as bright as Mizar as seen from Earth. There is nothing particularly special about Mizar. It is just a convenient reference star, which appears in the handle of the Big Dipper.

Star	Apparent Brightness relative to Mizar
Proxima Cen	2.54×10^{-4}
Cen A	6.72
Cen B	1.96
Barnard's Star	1.02×10^{-3}
Wolf 359	2.58×10^{-5}
BD+36 2147	6.66×10^{-3}
L726-8 A	6.56×10^{-5}
UV Cet B	4.12×10^{-5}
Sirius A	2.55×10^{1}
Sirius B	2.20×10^{-4}

EXERCISE 5-1: *Open a new Sheet in Excel by clicking the* **Sheet4** *tab. Create an Excel spreadsheet with the Star names (column A) and apparent brightnesses (column B) from the table above. Be sure to use the correct form for scientific notation in Excel! Calculate the logarithm of the apparent brightness for each of the stars and place it in column C.*

The **apparent visual magnitude _m_** of a star is computed from its apparent brightness by the formula:

$$m = -2.5 \times \log(\text{apparent brightness}) + 2.06 \tag{2}$$

where the constant 2.06 is needed to make these calculated magnitudes correspond to those derived by Hipparchus. The magnitude scale was not purposely designed by Hipparchus to be logarithmic; it just reflects the sensitivity of the human eye. The eye has evolved to be sensitive over a tremendous range of light levels, ranging from sunny days to a starlit night. To accomplish this remarkable feat it uses a shorthand of its own. Thus, a large range of sensitivity yields information that is compressed logarithmically and sent to our brains. Logarithmic scales are used in many other places, such as in describing sound levels (decibels) and earthquake vibration amplitudes (the Richter scale).

> **EXERCISE 5-2:** *Using your Excel spreadsheet, calculate the apparent magnitudes **m** for each of the stars using Equation 2 above and place the results in column D of your current spreadsheet.*

NOTE: Just like logarithms, astronomical magnitudes are <u>never</u> expressed in scientific notation. Always use standard notation. Also, just like logarithms, only the digits to the right of the decimal point are significant digits.

Lab Report Cover Page

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Fall 2022

Lab A
<u>Working with Numbers, Graphs, and the Computer</u>

Name: _____

Lab Section: _____

Date: _____

TABLE 1: EXERCISE 1-1

Convert the values given in the Standard Notation into Scientific Notation

Standard Notation	Scientific Notation
632,000,000,000	
632	
63.2	
632,000	
0.632	
0.000000632	
0.00632	
0.0000632	

Lab B
Energy Transport in Stars

(Adapted from *Contemporary Laboratory Experiences in Astronomy*)

PURPOSE:

To explore the nature of photons and how they interact with gas atoms and to understand how the absorption and emission lines we see in stellar spectra are produced as a result of these interactions. To understand why it takes a photon a long time to travel from the center to the surface of the Sun.

EQUIPMENT:

CLEA computer program ***Solar Energy Transfer***
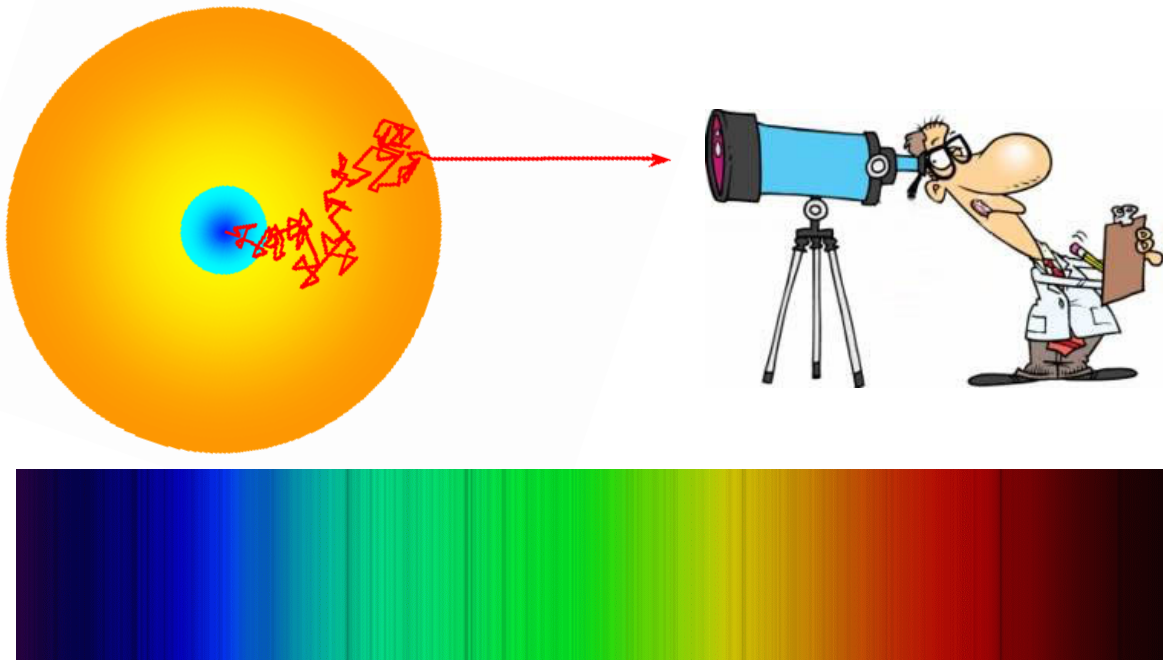
Figure1: TOP – *Schematic representation of a photon of light working its way from the center of the Sun to an observer's telescope.* **BOTTOM -** *The visible-light spectrum of the Sun showing dark absorption lines against a continuous background.*

Introduction

As you have learned (or will!) in your lecture class, the light we see coming from the Sun and the stars originates deep within their interiors in a region known as the *stellar core*. It is only here that the temperatures and densities are high enough to allow nuclear fusion reactions to occur. These reactions release huge amounts of energy, which provide the *thermal pressure* that helps a star withstand its own self-gravity and ultimately emerge from the star's surface as the light that makes the stars shine in the nighttime sky.

Much of the energy generated in a star's core of is carried to its surface (and beyond) by individual *photons* of light. As the photons travel outwards from the core, they continuously interact with the gas particles making up the mass of the star. For the purposes of this discussion, we can think of a star as being composed of two general regions:

- *The Interior* – the main body of the star. In this hot, dense region most atoms are fully *ionized*. I.e., their electrons have been stripped off, leaving behind a positively charged *ion*. The electrons, which can move independently through the gas, are called *free electrons*.

- *The Atmosphere* – the thin layer of gas that makes up the outermost skin of a star. In this cooler, lower density region, many electrons are attached to atomic nuclei, and no longer free to move independently through the gas. These are known as *bound electrons*.

There are several distinct types of interactions that may occur between photons and gas particles as the energy works its way to the stellar surface.

- *Electron Scattering* – which occurs when a photon encounters a free electron. The photon is absorbed by the electron and then immediately re-radiated in some random new direction (as if it had simply scattered off the electron). The photon may gain or lose some energy due to this interaction.

- *Free-Free Absorptio*n – in which a free electron in the vicinity of an atom or ion can absorb a passing photon, making the electron more energetic. The electron may later give up this extra energy at any time in the form of a new photon, again to be radiated in some random direction.

- *Bound-Free Absorption* – in which a photon is absorbed by an atom or ion, causing it to eject a (formerly) bound electron. This new free electron can *recombine* with another ionized atom some time later, giving rise to the release of a new photon in some random direction.

- *Bound-Bound Absorption* – in which an atom or ion absorbs a passing photon, causing one of its bound electrons to become more energetic (but still remain bound). The electron may later give up this extra energy in the form of a new photon, again to be radiated in some random direction.

Photons traveling through the hot dense stellar interior interact with the gas primarily through the first three processes (electron scattering, free-free absorption, and bound-free absorption). As a result, these photons travel a zig-zag path on their way out, as they are scattered or absorbed and reemitted in nearly random directions. See Figure 1. So many interactions occur that, in the Sun, it takes literally hundreds of thousands (to, perhaps, millions) of years for a typical photon to travel from the center to the surface!

The stellar atmosphere is cooler than the interior and largely transparent to most of the photons that reach it after traveling through the interior. Thus, these photons simply pass through the atmosphere and out into

space, creating a *continuous spectrum* and producing the light we observe from the stars. But some photons – with very specific energies – undergo the bound-bound absorption process and are prevented from leaving the stars. This produces gaps in the spectrum of the Sun and stars – at very specific energies – that are known as *absorption lines*. See the bottom part of Figure 1. Analysis of the emergent spectrum of light from stars, including the absorption lines, can provide a great deal of information on the physical conditions within a star.

Lab Procedure

The computer program that you will use in today's lab (***Solar Energy Transfer***) is designed to illustrate how energy flows from the energy-generating core of a star out into space. Specifically, you will see how individual bits of light, known as *photons*, travel from the core of the Sun to its surface, and how they interact with matter on their way into space. The program includes several simulations of how photons behave as they interact with atoms of gas.

To begin, click on the ***Solar Energy Transfer*** program in the "MSE Astrolab Student Folder" on your Desktop. Select **Log In...** from the menu and enter your name. Do not use punctuation marks. When the information is correct, click **OK** to continue.

EXERCISE 1: ENERGY FLOW THROUGH THE INTERIOR

In this first exercise, the "FLOW" simulation, you will observe how a photon works its way up from the center of the Sun (or other star) towards its surface. You will be shown a two-dimensional slice of the interior of a simulated star and observe how a photon diffuses outward from the core and how the numbers of layers of atoms used to model the simulated star affects the amount of time it takes for a photon to escape into space. As a result of the electron scattering, free-free, and bound-free processes, every time a photon interacts with matter, it is re-directed so that it travels in a new and nearly random direction. The resulting zig-zag path is called *random walk*. This is graphically demonstrated in the FLOW Simulation.

Step 1:   Start the program

Click on **Simulation**, then **Flow**, then **1 Photon** to start the energy FLOW simulation.

Step 2:   Perform a 5-layer simulation

Click ***Parameters, # of Layers…*** then type in "5" and click ***OK.*** Press ***Run Simulation*** and then observe the photon as it works its way out from the interior of the star to its surface and escapes into space. The simulation will end when the photon escapes and the number of times the photon interacted with an electron will be listed on the screen ("Interactions"). Record the number of interactions in the appropriate column of Table 1 (in the row labeled "Trial 1"). Repeat two more times (Trials 2 and 3) end enter the number of interactions into Table 1. After you have completed three trials, compute the mean number of interactions and enter it Table 1.

Note: Under ***Parameters, Trails*** you have the option of tracing out the photons' crooked path. Trails are visually interesting but remember that photons are not "eating" their way out of the star in any sense, they are simply bouncing around! Typically, the program is run with **Trails** set to **No.**

Step 3:   Perform 10, 15, 20, 30, and 40-layer simulations

Now repeat Step 2 using simulations containing 10, 15, 20, 30, and 40 layers of atoms. Perform 3 trials for each model and enter the results in Table 1.

Table 1: Flow Simulation

| *n* (number of layers) | Number of Interactions | | | | n^2 |
	Trial #1	Trial # 2	Trial #3	Average	
5					
10					
15					
20					
30					
40					

TIME SAVING HINT
You can open a second instance of *Solar Energy Transfer* and make the measurements for two (or more) different numbers of layers simultaneously.

Step 4:   Analyze the data

Using either the graphs provided at the end of the lab manual or Excel (as per your instructor's directions), make a plot of the number of interactions (y-axis) as a function of the number of layers in the simulation (x-axis). Sketch in a smooth curve that best fits your data. **If you are using Excel, use the *Trendline* feature and the *Power* option.** Be sure to show the equation of your Power fit on the plot.

It can be shown theoretically that the average number of interactions that occur before a photon can escape is very nearly n^2, where *n* is the number of layers in the model. Complete the last row of Table 1 to compare your results with this prediction.

> **Question 1:**   *How do the calculated values of n^2 compare with the average results from the simulations? If they are different, why? If you used Excel and fit your data with a "Power" Trendline, then the exponent in your Trendline equation will be 2 if your data agree exactly with the n^2 prediction. Is this the case? Are your results consistent with the prediction? (In answering this question note that there is a randomness to the scattering process. Scattered photons aren't aimed at any particular direction, so whether an atom happens to be in their way is a matter of chance. Think of it this way: If you flipped a coin 4 times, would you always expect to get 2 heads and 2 tails? What if you flipped it 40 times?)*

Write the average number of interactions needed for the 40-layer model on the classroom whiteboard. When the whole class has done this, average the results.

> **Question 2:** *What is the %-difference between the class average and 1600 (which is 40^2)? Use the following formula:*

$$\text{\%-difference} = \frac{\text{observed-expected}}{\text{expected}} \times 100\%$$

> *What is the %-difference between your 3-sample average and the expected value of 1600? Which result has the smallest %-difference (in terms of absolute value)? Is this what you should expect? Explain.*

You will have noticed that the time required for a photon to escape from the interior of the star got longer as the number of layers in the simulation increased. Luckily for you, we stopped at 40 layers. In a realistic model, the size of the layers would be comparable to the separation between atoms and many (MANY!) more than 40 layers would be required. However, since the lab period does not last for a million or so years, we limited the demonstration to 40 layers.

Click **Return** to end the FLOW Simulation

EXERCISE 2: A CLOSE-UP STUDY OF PHOTON/ATOM INTERACTIONS

Once a photon gets close to the surface of a star, and enters its atmosphere, it still has to escape into space. The atmosphere is essentially transparent to most photons and they easily pass through it. However, if the energy of a photon happens to match the "excitation energy" of an atom in the atmosphere, then the photon will be absorbed by the atom and later re-radiated in some random direction, i.e., the "bound-bound" process will occur. The excitation energy is the energy required for a bound electron in an atom to move to a higher energy state. Once the electron moves to the higher state, it immediately jumps back down to the lower state, releasing the photon in a random direction. In the simulations below, photons that match an atom's excitation energy are called "**line photons**." If a photon's energy is not properly matched to the atoms in a gas, the photon will pass through the gas largely unaffected. These are called "**continuum photons**."

Click on **Simulation** and choose **Interaction**.

This simulation allows you to study how individual photons interact with atoms. The display portrays a field of atoms in a gas. You can think of this as representing a stellar atmosphere. Photons will enter from the left, as if emerging from the stellar interior, and attempt to pass through the atmosphere and exit at the right. You can change the scale of the simulation by selecting **Large Scale** or **Close-Up** from the **View** menu. Try both! You will see the *electron cloud*s surrounding each atom. These are the regions where the *bound electrons* are found. The atomic *nuclei*, where the *protons* and *neutrons* reside, are buried deep in the center of each atom and are invisible at the scale of this simulation. These simplistic pictures are for reference only and are not drawn to scale. Atoms are seldom so round because electrons reside in complex levels above the nucleus called "shells," which are not necessarily spherical.

Select **Photon Type, Line** to choose this type of photon to be sent through the gassy region. **Line photons** have precisely the right amount of energy to be absorbed by the atoms in the gas. Absorbing a photon adds energy to the electrons in the atom, "exciting" them into higher energy levels. After a very brief moment, they fall to a lower energy state and release a photon in some random direction. The electrons are said to be undergoing *bound-bound transitions*, since they remain attached to the same atom even when excited. Click on the **Run** button to watch as line photons are sent from the left of the screen to pass through or

interact with the atoms. Click on the **Stop** button after 20 line photons are sent. Fill in Table 2.

Now repeat the process, but for continuum photons (***Photon Type, Continuum***). ***Continuum photons*** have the wrong energy to interact easily with the atoms, and often pass through them without incident. Occasionally, a continuum photon will scatter off of an electron, but not often. Click on the **Run** button to watch as continuum photons are sent from the left of the screen to pass through or interact with the atoms. Click on the **Stop** button after 20 continuum photons are sent. Fill in Table 2. You should have observed a significant difference in the number of interactions experienced by the two types of photons.

Table 2: Photon/Atom Interactions

	Line Photons	Continuum Photons
Number of photons sent		
Number of photons scattered		
Number of photons un-scattered		

Press ***Return*** to exit the INTERACTION simulation.

EXERCISE 3: EXPERIMENT MODE

Select **Experiment** from the **Simulation** menu to start.

In this final simulation, you will investigate the formation of an absorption line spectrum. You will shoot photons of a range of energies into a specific gas and determine the fraction which emerge unscathed and reach your detector. As in Exercise 2, you can imagine that this illustrates what happens as photons from a stellar interior attempt to pass through a stellar atmosphere. When few interactions occur (as in the case of a continuum photon) many emergent photons will be detected. When many interactions occur (as in the case of a line photon) few get detected because most are scattered away in other directions. A plot of the fraction of photons detected as a function of energy (or wavelength) is essentially a plot of the absorption spectrum of the gas and is the same as what you might observe for a star using a spectrograph attached to a real telescope. The energies or wavelengths you will be observing are in the "optical" spectral region, i.e., the region to which your eye is sensitive.

You have several different gases you can choose from for this study. They include thin gaseous clouds of calcium (Ca), hydrogen (H), magnesium (Mg), oxygen (O) and sodium (Na). Each kind of atom has its own characteristic set of energies and thus each of these gases will yield a different absorption spectrum. **You will perform your experiment using two different gases and compare the results**.

Step 1: Select the first gas.

Use the **Select Gas Atoms** option under **Parameters** to choose your first gas. Enter the name of the gas in the third column of Table 3 below.

Step 2: Select the number of photons

Use **# of Photons** from the **Parameters** menu to set the photon count at 25.

Step 3: Set the starting photon energy (i.e., wavelength)

Use **Change Photon Energy** from the **Parameters** menu to set the photon energy to 1.5 eV[1]. As you change the photon energy, the wavelength changes automatically since the two are related[2]. You can reset the counters and statistics using the **Reset** button.

Step 4: Run the Simulation for all photon energies

Run the simulation, sending 25 photons through the gas at each energy level from 1.5 through 3.2 eV. Use **Change Photon Energy** to adjust the energy after each 30-photon set. In Table 3, fill in the wavelength and the number of detected photons for each energy level as you proceed. For example, if you send 25 photons through the gas at the 2.3 eV energy level and 5 are detected, enter a 5 into the Table.

Step 5: Repeat the above three steps for a second gas, recording the results in Table 3.

> **TIME SAVING HINT**
> You can open a second instance of ***Solar Energy Transfer*** and make the measurements for your two gases simultaneously. Just be sure you don't mix up the numbers in Table 3!

After you have completed Table 3, make a separate graph of your results for each of the two gases. I.e., plot the number of photons detected (*y-axis*) vs. photon wavelength (*x-axis*). Use the graphs provided at the end of the lab manual or an Excel worksheet, as directed by your instructor. Draw a smooth curve through your data points. Be careful of especially deep absorption features. **If you are using Excel, you can have Excel join the data points with a smooth line.** Although this is a crude spectrum, you should be able to see one or more pronounced dips in the number of photons detected. The energy, and thus the wavelength, of the dips identify photon energies that easily energize the atom. Once energized, as you know from Exercise 2, the atoms will reemit the photons in random directions. Most of the photons will move in directions away from an observer viewing the atom head on, i.e., away from their original paths before the interaction.

> ***Question 3:*** *Look at the two plots you just made. Do you see any evidence for "lines" in the spectra? Explain how you recognize them. What energies and wavelengths do these correspond to? How do the two spectra compare? Can you tell them apart? Explain how the spectrum of a star might reveal the composition of its atmosphere.*

Click **Return** to exit the EXPERIMENT Simulation.

---

[1]An electron-volt is the amount of energy added to an electron when it is passed through an accelerating potential of one volt. It is a very small amount of energy.

[2]If you would like to calculate the wavelength (in meters!) by hand, the equation is:

$$\lambda = {}^{hc}/_{E}$$

Where: h = Plank's constant = 4.14 x 10^{-15} eV second; c = speed of light = 3 x 10^{8} meters/second; and E = energy of the photon in electron volts
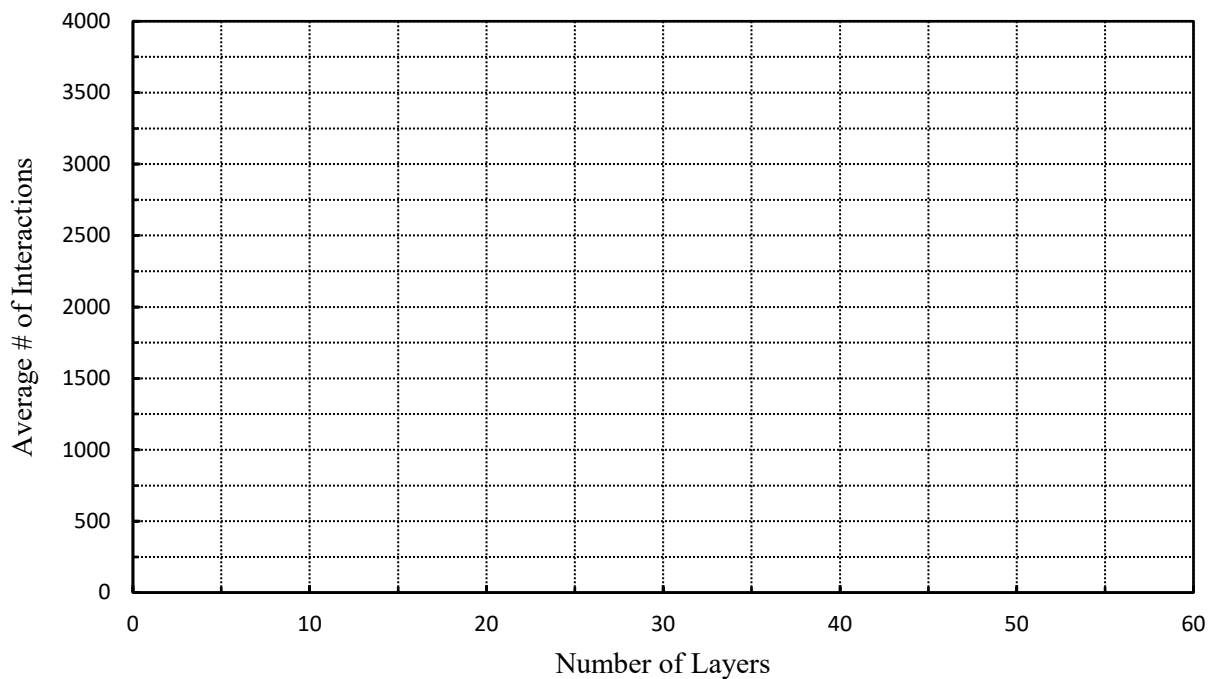
Table 3: Experiment Mode

Energy (eV)	Wavelength (nm)	Gas = Number Detected	Gas = Number Detected
1.5			
1.6			
1.7			
1.8			
1.9			
2.0			
2.1			
2.2			
2.3			
2.4			
2.5			
2.6			
2.7			
2.8			
2.9			
3.0			
3.1			
3.2			

Click **Quit** to exit the Solar Energy Transport program.

Summary Question

Question 4: *If this were a realistic attempt to simulate a photon escaping from the Sun, you would have to use a model with a <u>much</u> larger number of layers and the time you would have to wait for a photon to escape would be <u>much</u> longer. In fact, detailed calculations show that light generated in the center of our Sun can take up to 10,000,000 years to finally work its way to the surface! But the speed of light is 300,000 km per second. At that speed, light should only take 2 or 3 seconds to get out of the Sun (the Sun's radius is about 700,000 km). Using what you have learned in this lab, explain the apparent contradiction.*

EXERCISE 1: FLOW
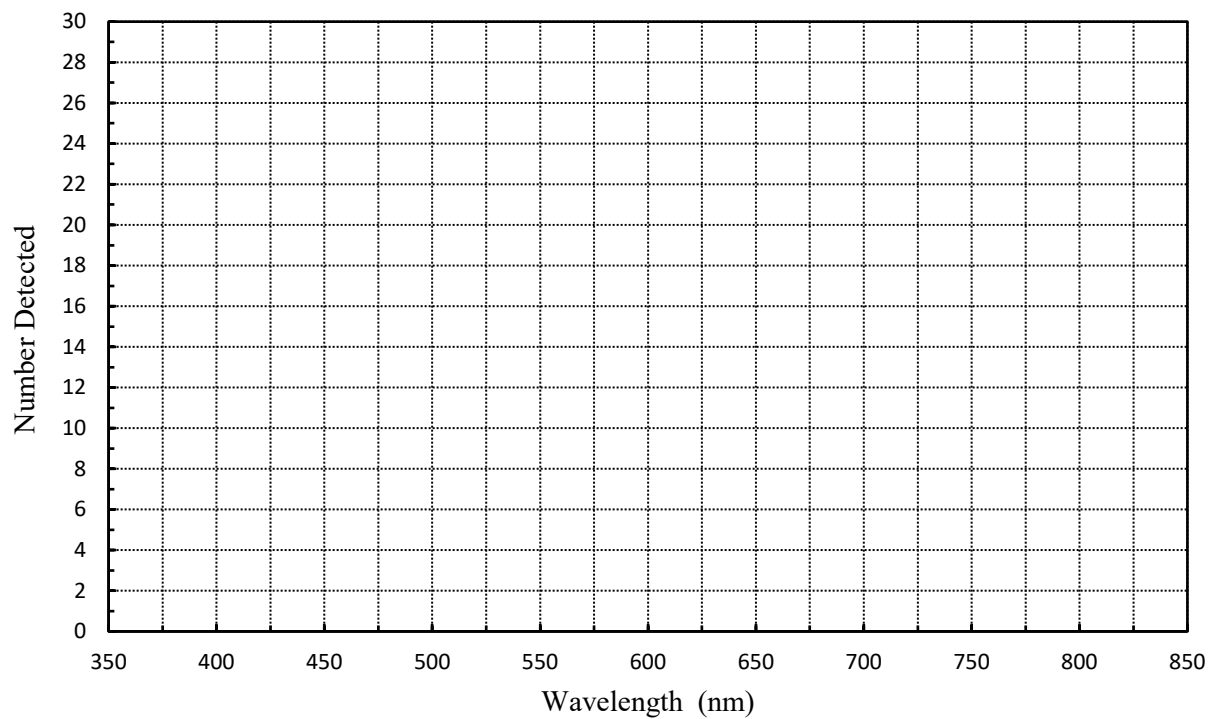
Average # of Interactions vs Number of Layers

EXERCISE 5: EXPERIMENT

Number Detected vs Wavelength (nm)

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Spring 2022

Lab B
<u>Energy Transport in Stars</u>

Name: _____

Lab Section: _____

Date: _____

Lab C
<u>Introduction to Spectroscopy</u>
(Adapted from *Laboratory Experiments in Astronomy*, Johnson & Canterna)

<u>PURPOSE:</u>

To examine the spectra of common light sources, learn how atomic spectra are related to atomic structure and see that the pattern of photons emitted by a hot gas can be used to identify the nature of the gas.

<u>EQUIPMENT:</u>

Transmission grating, meter stick, ruler, gas discharge lamps with power supplies, household light bulbs.
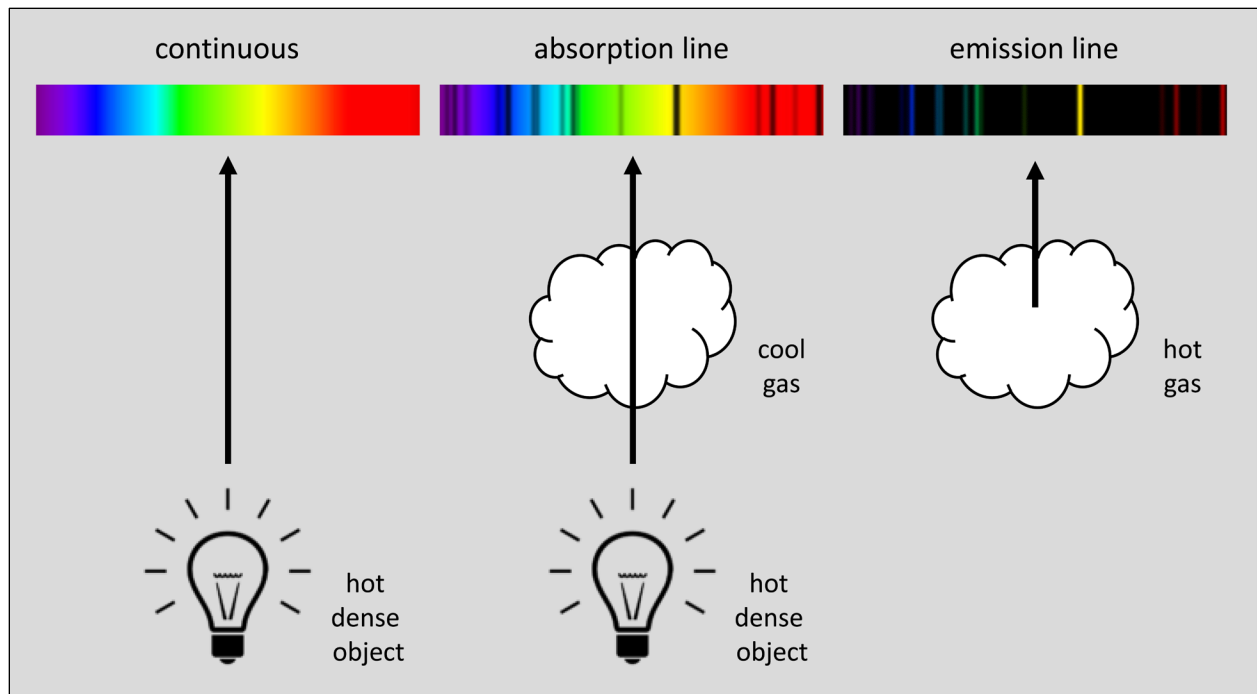
Figure 1: *The three types of spectra and how they form*

Introduction

Spectroscopy refers to the study of the electromagnetic radiation (i.e., light) emitted by matter, by dispersing the light into its component colors. The properties of this dispersed light, called a ***spectrum***, can reveal much about the properties of the emitting object. The three different types of spectra are illustrated in **Figure 1**. A ***continuous spectrum*** is produced by a hot dense object (e.g., the filament in a light bulb). In this case, light of all wavelengths is emitted. An ***absorption spectrum*** is produced when a continuous spectrum passes through a cloud of gas and some wavelengths are selectively absorbed by the gas. Finally, an ***emission spectrum*** arises when a low density gas is heated and selectively emits photons of specific wavelengths. The overall distribution of light in a continuous spectrum depends on the temperature of the emitting object, but not on its composition. For example, a 300˚K piece of metal emits the same continuous spectrum as a 300˚K human being. It is the emission and absorption spectra which can reveal the composition of an object, due to the atomic nature of matter.

An ***atom*** consists of two basic regions, a central ***nucleus*** (containing ***protons*** and ***neutrons***) and a surrounding ***electron cloud*** (containing ***electrons***). The positive electrical charges of the protons in a nucleus attract the negatively charged electrons in the cloud, holding an atom together. Due to the laws of quantum mechanics, the electrons are only allowed to occupy certain regions around the nucleus. The locations of these regions depend on the energy of the electrons: those with small amounts of energy must lie close to the nucleus, while those with more energy will lie farther away. These allowed regions are commonly referred to as ***energy levels.*** The energy level closest to the nucleus is called the ***ground state***. Electrons prefer to be in their ground states but, by the addition of energy, can move up to higher levels. In this case, the electron is said to be in an ***excited state***. Once there, the electrons will spontaneously jump back down towards the ground state, often stopping momentarily at intermediate levels on the way and shedding energy with each transition between levels. Each time an electron jumps between two levels, it must gain or lose energy equal to the energy difference between the two levels, i.e., $E_{upper} - E_{lower}$.

An absorption spectrum is formed when an atom absorbs a passing ***photon*** of light that has just the right amount of energy to allow an electron to jump from a lower level to a higher level. Most photons do not meet this requirement and will pass the atom without being absorbed. The missing photons will be revealed by their absence in the spectrum, i.e., by the presence of dark lines at specific wavelengths. An emission spectrum is formed when excited electrons drop from upper levels to lower levels, emitting photons with energy equal to the energy difference between the levels and producing a spectrum consisting of bright lines. The original excitation may have been caused by collisions with other particles in the hot gas. (See **Figure 1**.)

In both cases, the energy differences between the upper and lower levels are revealed by the photons' wavelengths λ. From the general relation between λ and the energy E of photons, we have:

$$\lambda = \frac{hc}{E_{photon}} = \frac{hc}{E_{upper} - E_{lower}} \tag{1}$$

where ***h*** and ***c*** are Planck's constant and the speed of light, respectively. Thus, the wavelengths measured for emission or absorption lines reveal the energy level differences of the atoms in a gas. Since these differences are unique for each type of atom, the pattern of observed lines can be used to identify the types of atoms present in the gas. For this reason, the emission and absorption lines are often thought of as the "fingerprints" of the atoms.

Lab Procedure

In this experiment, you will measure the wavelengths of photons emitted by hydrogen and by a "mystery gas," which you will attempt to identify by its unique spectral signature. The emission is produced by running an electrical current through a tube of gas (a ***discharge tube***). The energy supplied by the current causes the electrons in the individual gas atoms to jump from their ground states up into one of the excited states. The electrons then spontaneously jump back towards the ground state, losing the gained energy by emitting photons of light at specific wavelengths. The glow that you will see from the gas tube results from these emitted photons. The set of photons emitted from a glowing gas are referred to as the gas's ***emission lines***. Later in the lab, you will also examine the emergent spectrum from several common sources of household lighting.

EXERCISE 1:   The Hydrogen Spectrum

In the simplest of all atoms, the hydrogen atom, the nucleus consists of a single proton and the electron cloud contains a single electron. **Figure 2** schematically illustrates the hydrogen atom. The single proton is at the center and the single electron may occupy one of the discrete energy levels labeled $n = 1, 2, 3…$ in the Figure. The $n = 1$ level is hydrogen's ground state. Note that hydrogen has many energy levels and only the first six are shown in **Figure 2**.

The wavelengths shown in **Figure 2** (in units of *nanometers*, or nm) are those corresponding to the photons that would be emitted by an electron jumping between the two energy levels connected by the solid lines. Notice that jumps ending at the $n = 1$ level (called the "Lyman series") produce photons with ultraviolet wavelengths, and those ending at $n = 3$ ("Paschen series") produce infrared photons. Neither of these can be directly observed by the naked eye. The $n = 2$ level, however, is the "Goldilocks" level. Electron jumps ending at $n = 2$ produce photons in the optical spectral region, and thus are visible to the naked eye. These are known as the "Balmer series" or the "Balmer lines." There are many other series of transitions involving higher energy levels than shown in the figure. These involve photon energies increasingly far into the infrared region.

The first step in the lab will be to view and measure the wavelengths of three members of the Balmer series. The three emission lines you will see – red, blue-green, and violet – all come from electron jumps that end at the $n = 2$ energy level, as follows:
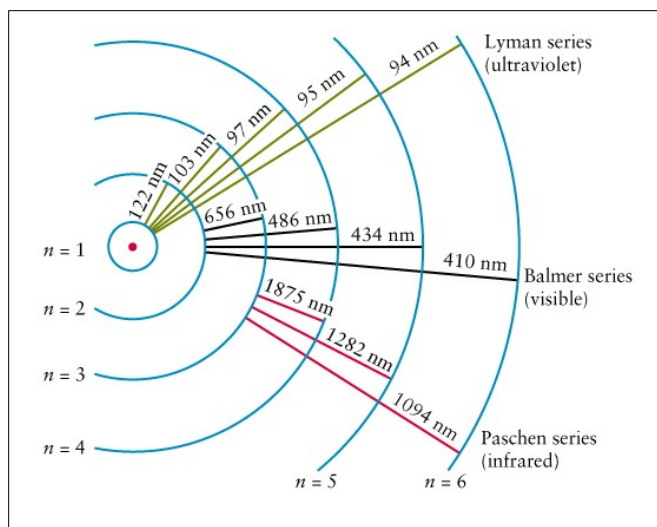

Figure 2:   *The energy levels in a hydrogen atom*

$n = 3 \rightarrow 2$:   [red]
$n = 4 \rightarrow 2$:   [blue-green]
$n = 5 \rightarrow 2$:   [violet]

Some of you may have really ultraviolet-sensitive eyes and see a fourth line, corresponding to

$n = 6 \rightarrow 2$:   [really violet]

You will accomplish the measurements by using a simple spectrometer, which is shown schematically in **Figure 3** below. With this setup, you will view the hydrogen discharge tube through a ***diffraction grating***, which is a piece of transparent glass or plastic with many parallel grooves etched closely together on its surface. The grating acts much like a prism, dispersing light into its component colors. As light from the discharge tube passes through the grating, its direction of travel is altered so that it appears to the eye to originate from a location displaced from the discharge tube. The size of the displacement depends on the wavelength of the emission line.

In your experimental setup, position the meter stick so that the hydrogen discharge tube is as close to its center. The emission lines will be seen projected against the meter stick. One spectrum will appear projected to the right of the center of the meter stick and an identical spectrum projected to the left. Even slight misalignment of the grating in your spectroscope can shift the spectrum to the left or right on the meter stick. To correct for this, you must average the distance *s* of a line in the left spectrum from the center of the meter stick with the corresponding distance in the right spectrum. ***The easiest way to do this is to record the number from the meter stick corresponding to the left position of a line and then from the right position.*** Record these values ***to the nearest millimeter*** in the top part of **Table 1**. The positive difference (or absolute value of the difference) of these numbers divided by two gives the best estimate of *s*. If you see two violet lines of hydrogen, measure the position of the brighter one.

Now measure the value of *x* to the nearest millimeter. As indicated in **Figure 3**, this is the distance between the front face of the meter stick and the transmission grating:

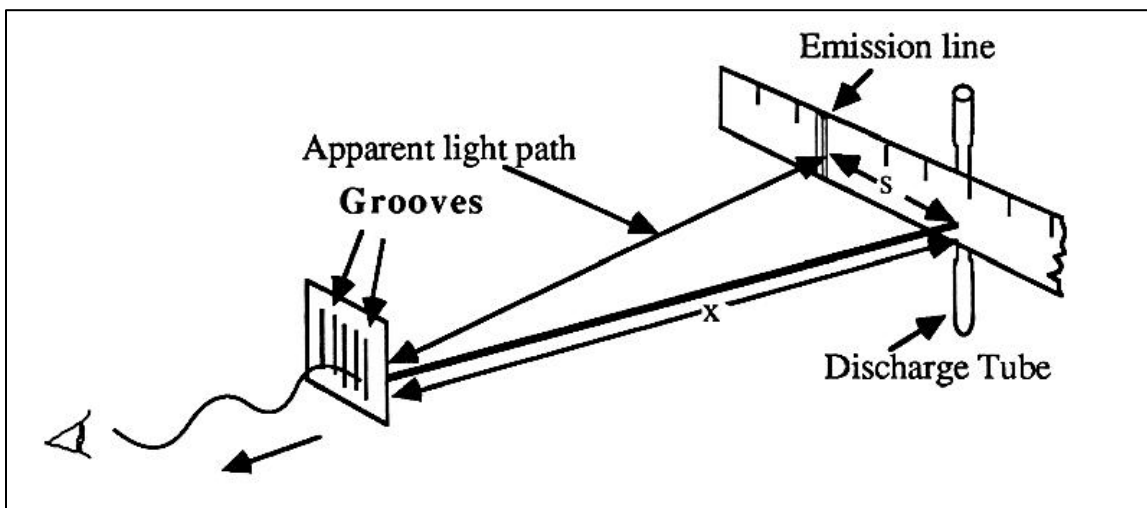$$x = \text{_____} \text{ cm}$$

Figure 3:   *The experimental setup*

C-4

The wavelength of the individual emission lines, λ, can now be determined from the following relation:

$$\lambda(\text{cm}) = \frac{s(\text{cm})}{g(\text{grooves/cm}) \times \sqrt{x^2(\text{cm}^2) + s^2(\text{cm}^2)}} \tag{3}$$

where x is the distance between the grating and the front face of the meter stick, s is the projected distance between the emission line and the discharge tube, and g is the number of grooves per centimeter on the diffraction grating. The wavelength λ will be in centimeters if x and s are in centimeters and g is in grooves/cm. Your diffraction grating has g = 5,000 grooves/cm.

Calculate the wavelength in centimeters for each emission line you observe using **Equation 3** and then convert these into units of Angstroms (1 Å = 10^{-8} cm). Record these data in the lower part of **Table 1**.

> **Question 1:** *How close were your measured wavelengths to the actual values? The red line (known as "Hα") has a wavelength of 6563 Å, the blue-green line ("Hβ") has λ = 4861 Å, and the violet line ("Hγ") has λ = 4341 Å. What are the absolute differences between your results and these true values? What are the percent errors? The percent error is given by:*

$$\%\text{-error} = \frac{\text{actual value - observed value}}{\text{actual value}} \times 100\% \tag{4}$$

> *Do you notice any sign of a "systematic error" in your results? (I.e., are all measurements too high or too low?) If so, how big is this systematic error? Any ideas on how this might arise?*

EXERCISE 2:   Identifying a "Mystery Gas"

Now switch your hydrogen discharge tube with a tube supplied by your instructor, containing a "Mystery Gas." (Labeled "A", "B", "C".) Be careful! The tubes are fragile and the hydrogen tube may be hot. If you are not comfortable switching the tubes, ask your instructor for assistance. Observe your Mystery Gas through the diffraction grating and fill in **Table 2**. Measure the wavelengths of as many individual lines as you can. (Add more lines to **Table 2** if you need to.) Make sure to list which Mystery Gas tube you have chosen to examine (i.e., "A", "B", or "C"). Compare your colors and wavelengths with the atomic spectrum wall chart in the lab or with **Figure 4** at the back of the lab.

> **Question 2:** *What kind of gas do you think your mystery tube is filled with? What features do you see that lead you to this conclusion? Are there any observed features not consistent with your selection?*

EXERCISE 3:   The Spectra of Common Light Sources

Go to the station at the back (or front) of the lab room, where your instructor has set up spectroscopes for you to observe the spectra of three common types of light bulbs: **incandescent**, **LED**, and **compact fluorescent** ("**CFL**"). Study the spectra carefully and answer the following questions:

> **Question 3:** *What does the spectrum of the **incandescent bulb** look like? Describe its properties. Which of the three types of spectra in Figure 1 most resemble this spectrum? What similarities and differences do you see between this spectrum and those you observed in Exercises 1 and 2? Does this spectrum tell you something about the physical process the bulb uses to generate light?*

> **Question 4:** *What does the spectrum of the **LED bulb** look like? Describe its properties.*

Which of the three types of spectra in Figure 1 most resemble this spectrum? What similarities and differences do you see between this spectrum and that of the incandescent bulb? Do you notice anything about the temperature of the LED bulb compared to the incandescent bulb? Does this tell you anything about the efficiency of each bulb?

Question 5: *What does the spectrum of the **CFL bulb** look like? Describe its properties. Which of the three types of spectra in Figure 1 most resemble this spectrum? What similarities and differences do you see between this spectrum and those you observed in Exercises 1 and 2 and for the other bulbs in this exercise? Does this spectrum tell you something about the physical process the bulb uses to generate light?*

Note, you may do a search on the internet to find out a little more about the types of bulbs you are looking at and the properties of their emitted spectra. One possible site can be found at www.comsol.com/blogs/calculating-the-emission-spectra-from-common-light-sources.

*** EXTRA CREDIT ***
Note: because this is extra credit, you may not ask your instructor for help!

EXERCISE 4:   The Size of an Atom

The size of an atom is determined by the size of its electron cloud. For most atoms this ranges in radius from about 0.3 to 3 Angstroms. As per the discussion in the Introduction, this size can vary for a given atom, depending on whether its electrons are in excited states or not. For hydrogen, the electron is usually found in the ground ($n = 1$) state, which has a radius of 0.53 Angstroms. This value, known as the "Bohr Radius" is usually considered to be the radius of a hydrogen atom.

Question 6: *So, in more understandable term, how big is a hydrogen atom? Assuming the Bohr Radius, how many hydrogen atoms would you have to line up side-by-side to span 1 inch? How many hydrogen atoms would it take to fill up the volume of a sugar cube, i.e., 1 cm³? Show all your work and explain your calculations.*

TABLE 1: Hydrogen

n	Line Color	Left Position (cm)	Right Position (cm)	s = (Right-Left)/2 (cm)
3	Red			
4	Blue-Green			
5	Violet			

n	Line Color	λ (cm)	λ (A)
3	Red		
4	Blue-Green		
5	Violet		

TABLE 2 : Mystery Gas _____

	Line Color	Left Position (cm)	Right Position (cm)	s = (Right-Left)/2 (cm)

	Line Color	λ (cm)	λ (A)

Spectrum Analysis
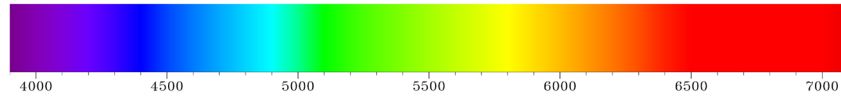
Hot solid objects emit light at all wavelengths, producing a "continuous" spectrum, as illustrated above for the wavelengths to which the human eye is sensitive. In contrast, hot gases emit light only at specific wavelengths, which are characteristic of each type of gas. Several examples of these "emission line" spectra are shown below.

Hydrogen

Helium

Carbon

Nitrogen

Oxygen

Neon

Argon

Krypton

Mercury
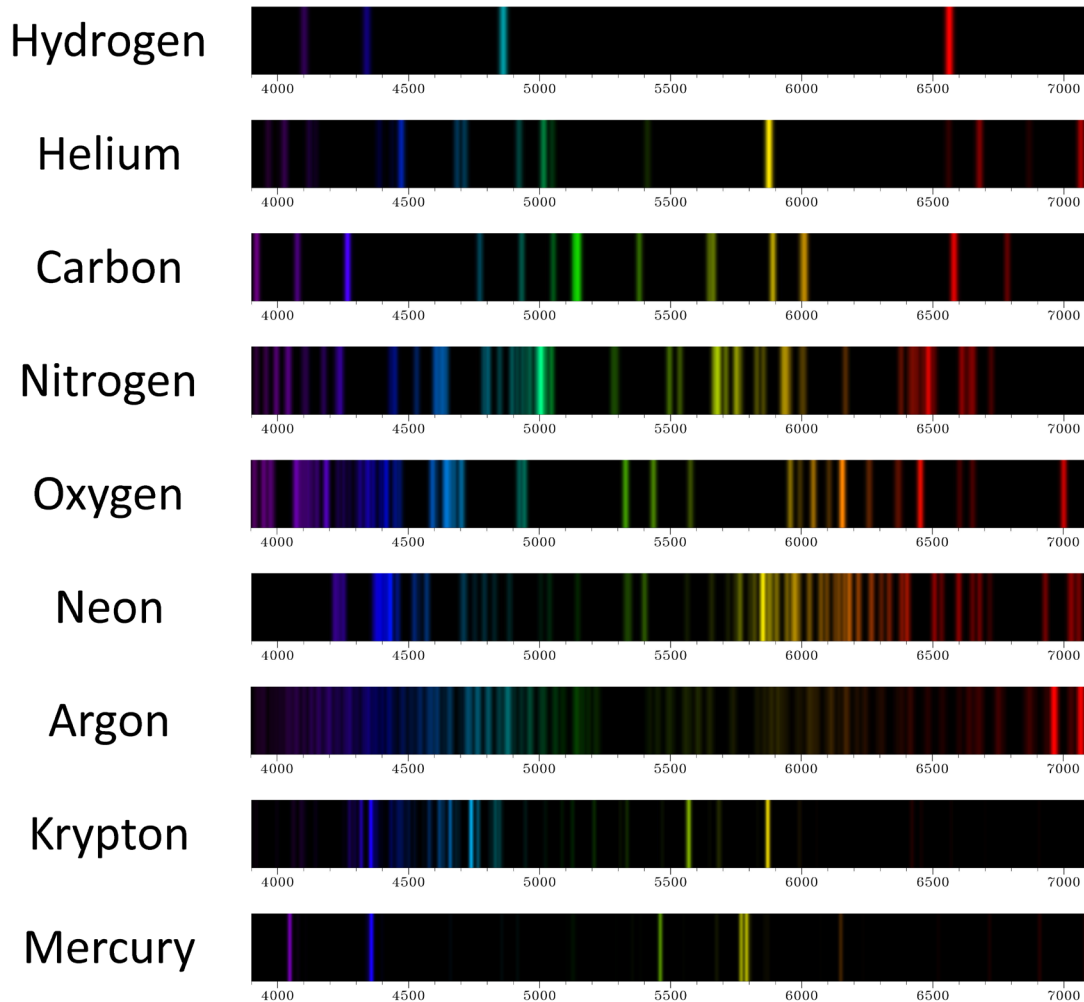
Figure 4: *Emission spectra of some common gases*

Lab Report Cover Page

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Spring 2022

Lab C
<u>Introduction to Spectroscopy</u>

Name: _____

Lab Section: _____

Date: _____

Lab D
Introduction to Optics

PURPOSE:

To study the ways in which light is refracted through lenses of various shapes, leading to an understanding of the principles behind the function of an astronomical telescope. To build and test a simple telescope using the Optics Bench.

MATERIALS:

Optics Bench kit (including lenses, mirrors and light sources), cotton gloves, ***Optics Bench*** simulation program, Microsoft Excel
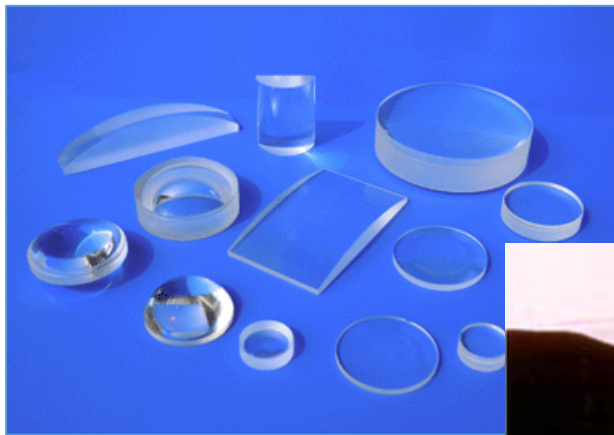
Figure 1: Clockwise from top left: a selection of lenses; Keck 10-meter reflecting telescope, image formation in the human eye.
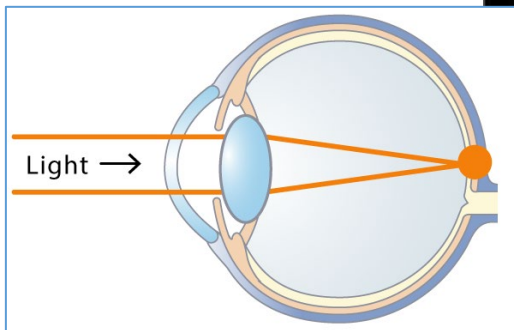
Introduction

**(NOTE: your instructor may ask you to answer introductory questions
1 thru 5 as a pre-lab exercise)**

SOME BACKGROUND:

In many ways, light can be thought of as a wave of energy propagating thru space (an "electromagnetic wave"). One of the properties of these waves is that they can change direction when they exit or enter a different medium as, for example, from air into glass or water. This property is called ***refraction***. The amount of refraction of the light wave depends on a property of the medium known as the ***index of refraction***. If you have ever looked into a fish tank or – as illustrated in **Figure 2** – at a straw in a glass of water, you have probably noticed the strange affects that arise from this phenomenon.
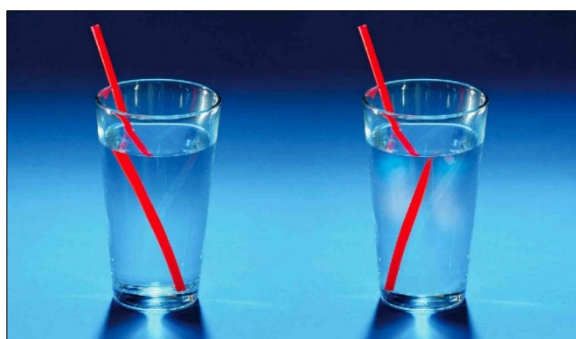
Figure 2: *Refraction*

Now let's look at what happens when light passes through a ***convex lens***. This is a smooth, transparent, circular piece of glass which is thicker in the middle than around the edges, as shown in the top half of **Figure 3**. Light from a very distant point source (which could be a star) arrives at this lens from the left as a bundle of light rays parallel to the ***principal axis*** of the lens (dotted line). The rays are refracted as they pass through the two surfaces of the lens and intersect on the right at point ***F***, known as the ***focus*** or ***focal point***, which is located a distance *f* from the lens center. This behavior gives rise to the general name for this type of lens, namely, a ***converging lens***. The distance *f* is called the ***focal length*** and is often used to characterize a lens. It is usually measured in meters, centimeters, or millimeters. Importantly, an image of the distant object, called a ***real image*** will be formed at point ***F***. That is, if a screen (or a piece of paper) were placed at point ***F***, an image of the distant source would appear projected on the screen

For astronomical uses, the important characteristic of a converging lens is its ability to gather and concentrate light into a small place. This is governed by the physical size of the lens. It is pretty straightforward to imagine that the larger the lens, the more light rays it will intercept. The more light rays a lens can intercept and concentrate in one place, the brighter the image will be. Thus, bigger lenses can form detectable images of fainter objects, such as stars or distant galaxies. In fact, the amount of light a lens can intercept is related to the *area* of the lens, which is proportional to the radius R of the lens squared. (Recall that the area of a circle is given by $A = \pi R^2$.)

A different sort of behavior arises from a ***concave lens***, which is a lens that is thicker around the edges than at the center, as shown in the lower half of **Figure 3**. This shape causes an incoming beam of parallel light rays to spread out and, as a result, this type of lens is known as a ***diverging lens***. Due to the divergence, the
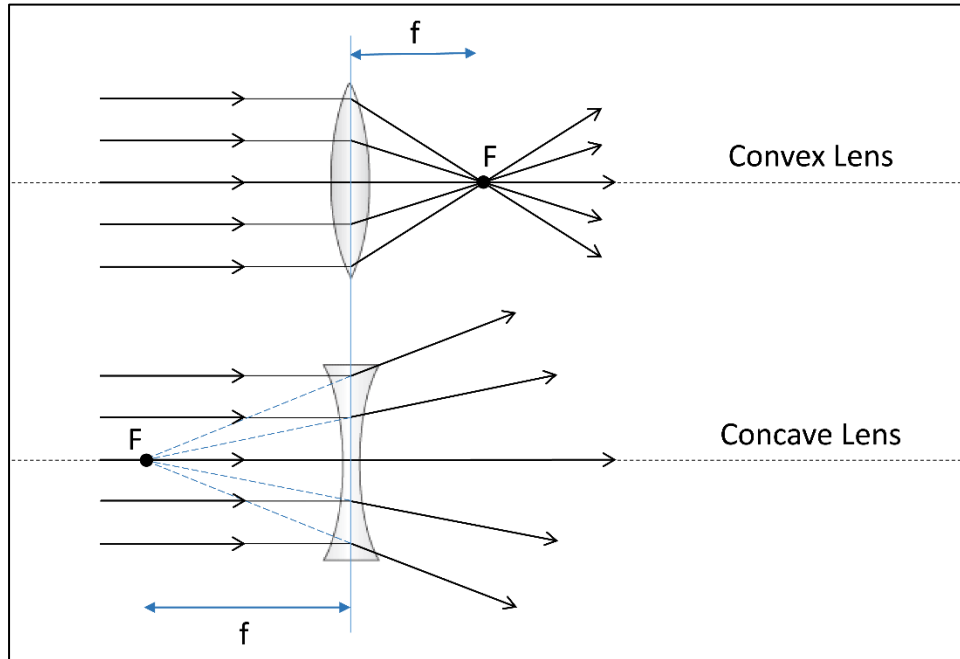
Figure 3: *Refraction through convex and concave lenses*

light rays *appear* to emanate from a position at the focal point ***F*** on the same side of the lens as the distant object. Concave lenses do not form real images. I.e., a screen placed at point ***F*** will do nothing except block the incoming beam of light!

IMAGE FORMATION WITH CONVEX LENSES:

We'll now consider objects that are not distant point sources, but rather have a larger angular size, i.e., ***extended objects***. Light from a point on such an object spreads out, striking a convex lens as seen in **Figure 4** below:

The sample light rays 1, 2, and 3 that travel from the tip of the arrow through the lens all take different paths. Ray 1 starts out parallel to the principal axis of the lens (dotted line) and is refracted through the focus on the right side of the lens. Ray 2, which happens to travel through the center of the lens, is not refracted at all and continues until it crosses the path of Ray 1. Ray 3 travels in such a way that it passes through the focus on the left side of the lens and is refracted into a ray parallel to the axis of the lens and intercepts Rays 1 and 2. If we trace rays from other parts of the object through the lens we will find that the lens forms an ***inverted real image*** of the object at d_i. Once again, this is known as a "real" image because, if a piece of paper or a CCD image sensor (such as in your camera or cell phone) were placed at this point, an image of the object would be seen or recorded. It can be shown – using basic geometry – that the location of a real image is given by:

$$\frac{1}{d_i} = \frac{1}{f} - \frac{1}{d_o}$$

(2)

where d_o and d_i are the object and image distances, respectively, and f is the focal length of the lens. If an object is extremely far away (i.e., large d_o), then the image forms at the focal point. The human eye functions exactly as shown in **Figure 4**. The lens of our eye forms a real image on the back of our eyeball (the retina)
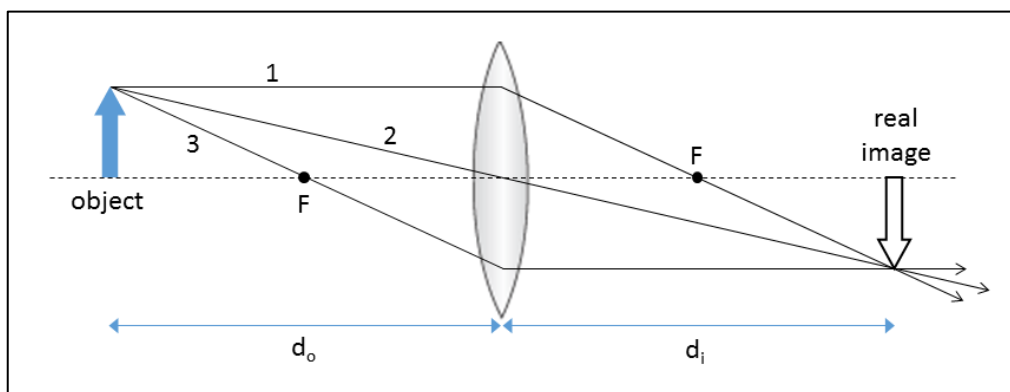
Figure 4: Image formation with a convex lens

and, yes, it *is* upside-down. A healthy eye is able to focus on both nearby and distant objects by adjusting its focal length so that the image always forms on the retina. (See **Figure 1** on the first page of the lab.)

THE REFRACTING TELESCOPE:

A ***refracting telescope*** uses two (or more) lenses arranged as shown in **Figure 5** below. These telescopes utilize the ability of lenses to both concentrate light from a larger area into a smaller area and to provide magnification of the resultant image. Essentially, the larger ***objective lens*** gathers a lot of light from the source, producing a bright real image, and then the ***eyepiece*** lens directs a magnified version of this image into our eyes (or a camera). The eyepiece is positioned so that its focus coincides with position of the real image formed by the objective lens. **Figure 5** shows the case for a very distant object, for which the real image is located at the objective's focal point. The size of the objective lens, which is always convex, determines a telescope's ***light gathering power***. Telescope eyepieces can consist of either convex lenses or concave lenses, as shown in the top and bottom portions of **Figure 5**, respectively. Notice that the convex and concave eyepieces are located on opposite sides of the objective's real image. The chief difference in the final image produced by the telescope is whether it is inverted (convex eyepiece) or upright (concave eyepiece).

It can be shown that the magnification of these simple telescopes is given by the ratio of the objective lens to eyepiece lens focal lengths, regardless of whether the eyepiece is convex or concave:

$$m = \frac{f_o}{f_e} \tag{3}$$

Telescopes usually come equipped with several eyepieces, of differing focal lengths, providing different levels of magnification.

INTRODUCTORY QUESTIONS:

> ***Question 1:*** *About how long do you think the focal length of the lens in a human eye is? How did you arrive at this estimate?*
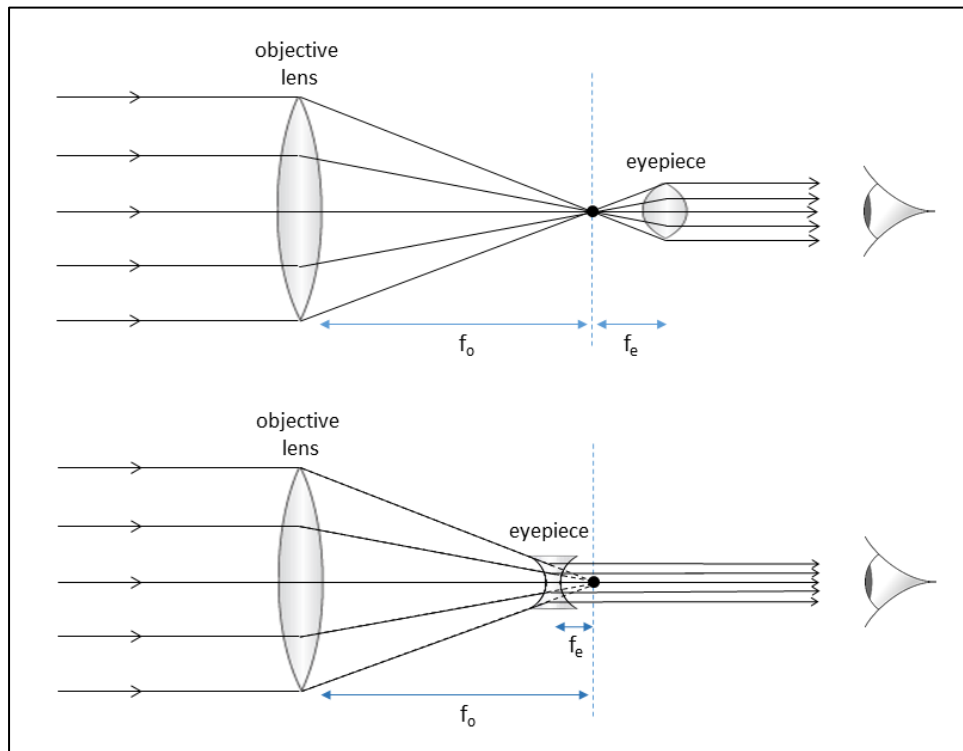
D-4

***Figure** 5: Simple telescopes utilizing convex (top) and concave (bottom) eyepieces*

Question 2: *Why do we see objects right side up if the image on our retina is upside down?*

Question 3: *In the days before smart phones and the internet, children used to amuse themselves outdoors on sunny days using single convex lenses to make small fires. Given what you have learned above, how is this possible?*

Question 4: *The Villanova University Public Observatory, which many of you will use this semester, has an objective element with a diameter of 14" and a focal length of f = 3900mm. The focal lengths of the eyepieces on this telescope are 10mm, 15mm, 23mm, and 32mm. What magnification is achieved with each of these eyepieces? Explain your results. I.e., what equations and inputs did you use?*

Lab Procedure

In this lab, you will investigate the properties of lenses and build a refracting telescope using both a computer simulation and a set of real lenses.

I. *Working with a Simulation*

First let's examine the properties of lenses using a simulation of an optical bench. Begin by starting the **Optics Bench** program found in the "MSE Astrolab" folder on your computer. Be sure that this program opens in the *Internet Explorer* browser and set the browser's "zoom" to 200%. You will be presented with the following screen:

Screenshot 1

The yellow line through the middle is the *optical axis* (or *principal axis*) of our system. The buttons on the top allow you to pick various types of sources and the buttons on the bottom are for selection of various optical elements. The yellow rectangle on the bottom left indicates the *x* and *y* positions of the cursor on the screen.

The first step is to add a convex lens to the simulation:

- Put the cursor on the **Lens** box and left click the mouse. The box will depress but nothing much else will happen.
- Move the cursor to the yellow line and left-click mouse. A convex lens symbol will appear. The position of the lens along the principal axis (*x*) and the lens' focal length (*fl*) will be displayed below the lens.
- Move the cursor over the lens until the cursor shows a hand symbol. Then left-click the mouse and hold while dragging the lens to a point near the middle of the screen. The *x* position will be about 5. The screen will look something like this:
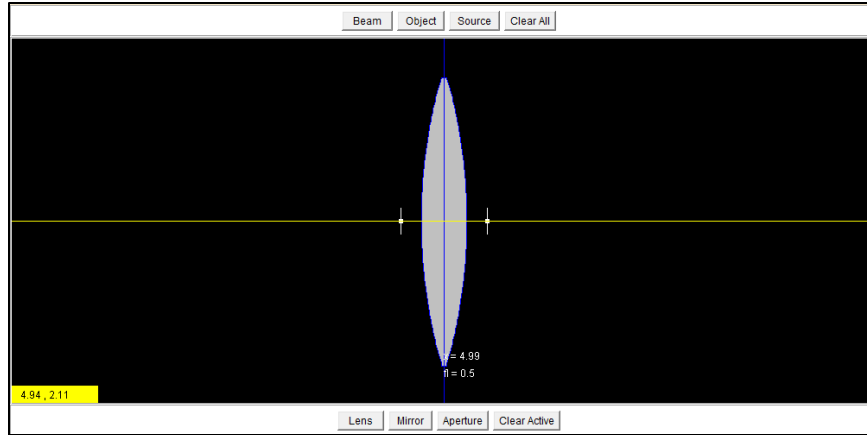
Screenshot 2

- Change the focal length of the lens by clicking and dragging either one of the two short vertical white lines. After experimenting with this, set the value to about *fl = 1.0* (it doesn't have to be exact).

    Question 5: *Imagine two convex lenses of the same diameter and the same composition. One is much thicker in the middle than the other. Which lens will have the shorter focal length?*

Now add a distant light source and observe the effect of refraction through the convex lens:

- Select the **Beam** box and place the beam source near the left edge of the screen. These parallel rays simulate the light coming from a very distant source, as in astronomical applications.

Notice where the parallel rays converge. Try moving the beam source around, noting the location of the convergent point. Try changing the focal length of the lens and, again, moving the beam source around and noting the location of the convergent point. An inverted **real image** of the distant source is formed at this location. I.e., a screen or a CCD detector placed here would record an image of the source.

    Question 6: *Where is the convergent point with respect to the lens' focal point? How does the convergent point change when the beam source is moved around? How does it change when the focal length of the lens is changes?*

Next add a nearby light source and observe the image formation:

- Left-click the beam source to activate it and then click the ***Clear Active*** box. This will erase the beam, leaving only the lens on the screen. Return the focal length of the lens to 1.0.
- Select the ***Object*** box and place the object (a vertical arrow) at approximately *x = 3.0*. Set the height of the object to about *h = 0.5*. The screen should look something like this:
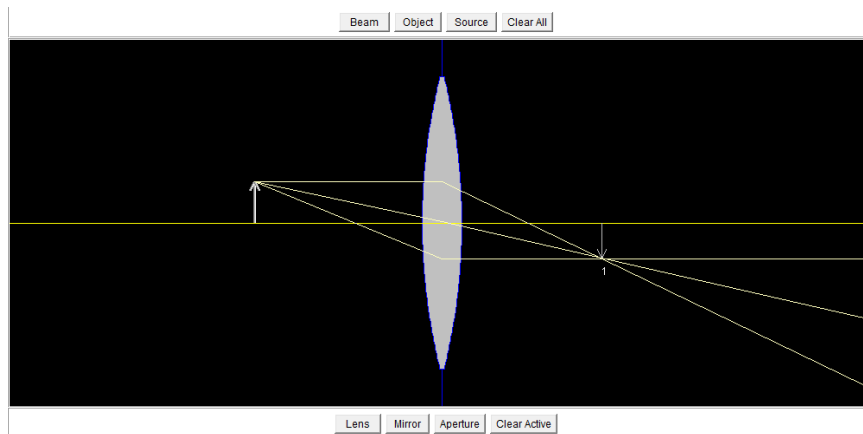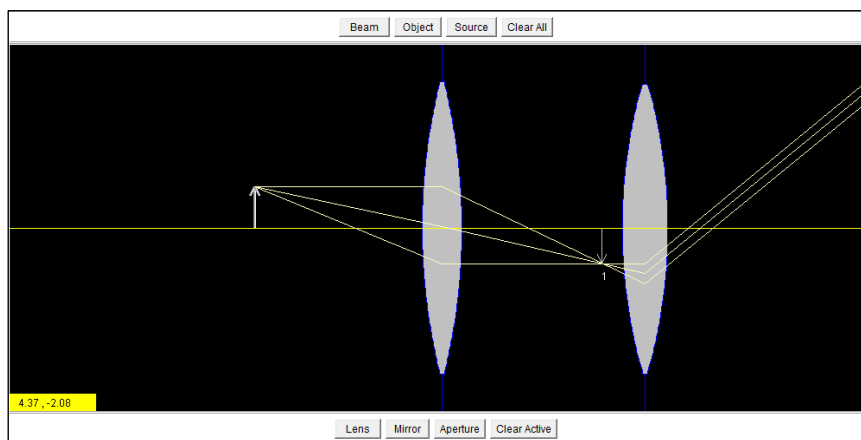
Screenshot 3

The light rays from a nearby source do not arrive at the lens parallel and, as a result, the ***real image*** formed by the lens (labeled "1" in the simulation) is not located at the focal point of the lens. (This situation is analogous to that shown in **Figure 4** of the Introduction.) Now move the source around and observe how the location of the real image changes with respect to the focal point of the lens.

> ***Question 7:*** *When you move the object closer to the lens, how does the location of the real image change? When you move it farther away? If you could move the object really far away – off the left side of the screen – where do you think the real image would be located? Is this consistent with expectations from Equation 2? Explain.*

Now it is time to build a telescope. To do this, we must add a second lens to the simulation.

- Click the **Lens** box and place the new lens such that its focal point lies exactly on the real image ("1") formed by the first lens. This lens has a focal length of 0.5. The simulation should now look like this:

Screenshot 4

As noted in the Introduction, the first lens in the simulation (the one closest to the object) is called the ***objective lens***, and the second lens is called the ***eyepiece***. The objective lens always has a longer focal

length than the eyepiece. Notice that the rays exiting the eyepiece are parallel to each other. This occurs when the telescope is perfectly focused. These parallel rays will be reconstructed by a normal human eye into a sharp image of the light source. The image will appear magnified by the ratio of focal lengths, as shown in **Equation 3**.

When an object is relatively near, as in terrestrial applications, the separation between the objective and the eyepiece lenses depends on the object distance. To see this, perform the following experiment:

- Note the current separation between the objective and eyepiece (i.e., record the *x-positions* and compute the difference).
- Move the object farther from the objective (but not all the way to the left of the screen), refocus the telescope (i.e., make the rays exiting the eyepiece parallel) and measure the separation between the objective and eyepiece. Note the location of the eyepiece's focal point with respect to the real image formed by the objective ("1").
- Move the object all the way to the left of the screen and repeat the step above.

    Question 8: After you first moved the object, was the telescope till in focus? I.e., were the rays exiting the eyepiece still parallel? After you refocused, how did the new separation between objective and eyepiece compare to the original value? After the second move of the object (all the way to the left), did you need to refocus the telescope again? How does the new separation between objective and eyepiece compare with the previous values?

    Question 9: In general, where must the eyepiece be located, with respect to the real image formed by the objective, for a telescope to be focused? How does this relate to what you are doing when you turn the focus wheel of, say, a pair of binoculars? Once an astronomical telescope is focused, should it ever need to be refocused? Explain. (Hint, an "astronomical" telescope is only used to observe extremely distant objects. For the purposes of Equation 2, the distances can be considered infinite.)

II. *Working With Real Lenses:*

OK, now let's try the above experiments using a real Optical Bench and actual lenses. ***Please note that you should never touch the surfaces of the lenses with your bare fingers. Use lens paper or cotton gloves, if available, or only handle the lenses by their edges.***

MEASURING FOCAL LENGTHS:

- Mount the light source box at one end of the optical bench, so that it is located at 0.0 cm according to the metric scale on the side of the optical bench. Plug it in. An illuminated target should be visible.
- Mount the white screen at the other end of the optical bench (the exact position is not important, since you'll be moving the screen).

You will now use the convex lenses to form real images and test Equation 2.

- Remove the 250 mm focal length convex lens from the case and mount it at a position of 60.0 cm according to the metric scale of the optical bench.
- Move the white screen back and forth until a sharp image of the light source appear on it. Do this carefully to achieve the sharpest possible image. This setup is analogous to that shown in **Figure 4**

and you are looking at the ***real image*** formed on the screen by the 250 mm convex lens! Note the size and orientation of the image (i.e., rightside-up or inverted).

- Measure the values of d_o and d_e (see **Figure 4**) and enter them, along with the image orientation, in Table 1.
- Repeat the three steps above for the 200 mm, 100 mm, and 50 mm convex lenses, always placing the lenses at the 60.0 cm point of the optical bench. Each time, note the properties of the real images formed by the convex lenses.

Table 1

Optical Element	d_o (cm)	d_i (cm)	Image Orientation?	f (cm)
250 mm convex lens				
200 mm convex lens				
100 mm convex lens				
50 mm convex lens				

Now use **Equation 2** to compute the focal length inferred for each lens by your measurements of d_o and d_e. Enter the results in the last column of Table 1.

> **Question 10:** *How did you do? I.e., how did your computed focal lengths compare with the expected values (i.e., 250 mm, 200 mm, etc.)? Can we trust Equation 2 to predict how a real image is formed by a convex lens? Does the size of the real image produced by a convex lens depend on the focal length of the lens? If so, how?*

Try the same experiment with a **concave** lens:

- Remove whichever convex lens is currently on your optical bench and insert the 150 mm **concave** lens at a position of 60.0 cm according to the metric scale of the optical bench.
- Carefully move the screen around and try to form a real image on it.

> **Question 11:** *Were you able to form a sharp real image on the screen using the concave lens? Is this consistent with discussion in the lab Introduction? How?*

BUILDING A TELESCOPE WITH A CONVEX EYEPIECE:

Now it's time to build an actual refracting telescope using your lens kit. You will use the 250 mm lens as the objective and the 100 mm convex lens as the eyepiece

- Place the 250 mm convex lens on the optics bench at a position of 60.0 cm.

The eyepiece (the 100 mm convex lens) must be placed at a position such that the telescope will be in focus. As you learned above, this lens should be located so that its focal point coincides with the real image formed from the objective lens (as in **Figure 5** or **Screenshot 4**). As a result, the distance between the eyepiece lens and the objective lens is:

$$d(\text{eyepiece-to-objective}) = d_i + f_{eyepiece} \qquad (4)$$

Compute this distance using the value of d_i that you already measured for the 250 mm lens (in **Table 1**) and the focal length f of the 100 mm eyepiece lens (i.e., 100 mm). Record this predicted value in the first row of **Table 2.**

Table 2

Telescope with 100 mm convex eyepiece	
d(eyepiece-to-objective)predicted	cm
d(eyepiece-to-objective)observed	cm
Telescope with 150 mm concave eyepiece	
d(eyepiece-to-objective)predicted	cm
d(eyepiece-to-objective)observed	cm

- Place the 100 mm convex eyepiece lens on the optical bench such that its distance from the objective is the value you just entered in **Table 2**.

Observe the lighted target through the eyepiece. Adjust the focus by moving the eyepiece, if necessary, to get the clearest possible image. If you move the lens, record the new eyepiece-to-objective distance in the second row of **Table 2.**

> **Question 12:** *Describe the image you see through the eyepiece. Is the image sharp or blurry? (Note: if you answered "blurry," go back and fine-tune the location of the eyepiece!) What is its orientation? What is its theoretical magnification (from Equation 3)? Is this consistent with an "eyeball" estimate of the magnification? Did you need to adjust the focus of the telescope? If so, what reasons can you think of that may have led to the predicted focus being less than optimal?*

BUILDING A TELESCOPE WITH A CONCAVE EYEPIECE:

As a final exercise, you will replace the convex eyepiece with a concave one. In this case, the predicted separation between the eyepiece and objective lenses is given by:

$$d(\text{eyepiece-to-objective}) = d_i - f_{eyepiece} \qquad (5)$$

The difference between **Equations 4** and **5** arises because, as can be seen in **Figure 5**, a concave eyepiece must be placed *between* an objective lens and its real image. Compute this predicted distance using the value of d_i that you already measured for the 250 mm lens (in **Table 1**) and the focal length f of the 150 mm concave eyepiece lens (i.e., 150 mm). Record this predicted value in the bottom half of **Table 2.**

- Remove the convex eyepiece from your telescope.
- Place the 150 mm concave eyepiece lens on the optical bench such that its distance from the objective is the value you just entered in **Table 2**.

Once again, observe the lighted target through the eyepiece. Adjust the focus by moving the eyepiece, if necessary, to get the clearest possible image. If you move the lens, record the new eyepiece-to-objective distance in the bottom row of **Table 2.**

> ***Question 13:*** *Describe the image you see through the eyepiece. Is the image sharp or blurry? (Note: if you answered "blurry", go back and fine-tune the location of the eyepiece!) What is its orientation? What is its theoretical magnification (from Equation 3)? Is this consistent with an "eyeball" estimate of the magnification? Did you need to adjust the focus of the telescope? If so, what reasons can you think of that may have led to the predicted focus being less than optimal?*

Before leaving the lab please do the following

1. Put all components of the optical kit back in the box in the original locations and let the instructor verify that all parts are there.
2. Place the white gloves back in the proper container (small or large) on the front desk.
3. Turn in your lab report.

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Spring 2022

Lab D
<u>Introduction to Optics</u>

Name: _____

Lab Section: _____

Date: _____

Lab E
Building a Galileoscope

PURPOSE:

To construct a replica of one of the first telescopes used by the Italian physicist/astronomer Galileo Galilei in the early 1600's. To make observations with the telescope and characterize the properties of the lenses and the equality of the images.

MATERIALS:

"Galileoscope" telescope kit.

Clockwise from top: Galileo's 20-power refracting telescope, Keck 10-meter reflecting telescope, modern small 'go to' telescope.

The Telescope

Sometime in the early 1600s, so the story goes, some Dutch children noticed that if you looked through two convex lenses held in line, you could see distant objects more clearly. History, however, credits the Dutch opticians Hans Lippershey, Zacharias Janssen, and Jacob Metius with purposely making the first telescopes. They applied for a patent for the device in 1609, but were refused because the device was so easy to duplicate. Lippershey noted that "counting coins from afar" was a suitable use for the invention! A year later, the Italian scientist Galileo made many improvements to the telescope design and was one of the first to use it to systematically study the night sky. Most telescopes of the time were used to observe terrestrial objects, such as ships popping up over the horizon, or as amusement. Low power "telescopes" were also used as opera glasses. These telescopes utilized a large "objective" lens and a smaller "eyepiece" lens and are known as ***refracting telescopes***.

As technology improved, it became possible to make bigger and bigger lenses for telescope use. As the lens size increased, however, it became clear that there was a limit to how large a practical telescope lens could be. The reason is that, above a certain size, the lens actually sags under its own weight and is no longer able to hold its shape under the force of gravity. This size turns out to be about 40 inches in diameter. Another problem with lenses is that they have two sides that need to be polished to the correct shape to form an image.

Early in the history of optics making it was learned that a curved polished or silvered mirror could, like a lens, form an image. Anyone who has looked at a silvered Christmas ornament has probably seen this effect. The use of mirrors as the main optical element of a telescope has a number of advantage over the use of lenses. Notably, using modern technology, very large, lightweight mirrors can be produced whose precisely-ground curved shapes can be maintained through the use of actuators mounted on the backs of the mirrors. These kinds of telescopes are known as ***reflecting telescopes***. All large modern telescopes are of the reflecting design.

Astronomers use telescopes to collect enough light to allow study of astronomical objects which are much fainter that the human eye can see. For this reason, telescopes are sometimes called "light buckets."

> ***Question 1:*** *Consider that the human eye has a maximum pupil opening of only about 7mm, compared to even a small telescope whose objective (front) lens might have a diameter of 50mm or more. Which do you think could collect more light from a star? Why?*

> ***Question 2:*** *Is the amount of light collected <u>directly proportional</u> to the diameter of the instrument (i.e., eye or telescope), or to some other property? Explain.*

The Galileo Telescope

The telescope you will be building in lab today is a 2009 *Galileoscope*™, which was created to help celebrate the 400th anniversary (in 2009) of the Italian physicist/astronomer Galileo's first use of the telescope to study the sky. The *Galileoscope* is a model of the type of telescope used by Galileo and other Renaissance astronomers in the early 17th century, but with modern high-quality lenses (which are vastly better than the primitive optics of the 17th century).   You can read more about Galileo's telescopes and discoveries at http://www.skyandtelescope.com/astronomy-resources/stargazing-with-galileo/

This telescope is a two lens system, consisting of a large objective lens and a smaller eyepiece. The main difference between Galileo's telescope and the simple telescope you built in the previous lab is that the Galileo design uses a concave lens for the eyepiece, as shown in the figure below. The advantage of this design is that a concave eyepiece keeps the image right side up and is more convenient for terrestrial viewing, which was the main use of telescopes in Galileo's time.
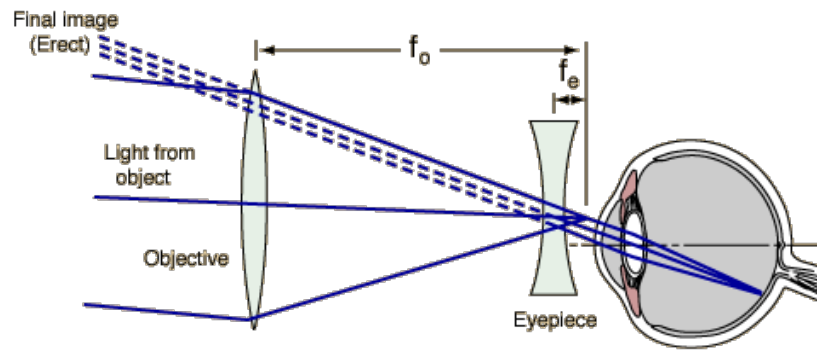
Figure 1: Image formation in Galileo's telescope.

Notice the placement of the eyepiece relative to the focus of the objective lens compared with the simple telescope you constructed in the last lab. (The focus of the objective lens is behind the eyepiece lens.)

It can be shown that the magnification produced by these simple telescopes is given by the ratio of the objective lens focal length to the eyepiece lens focal length:
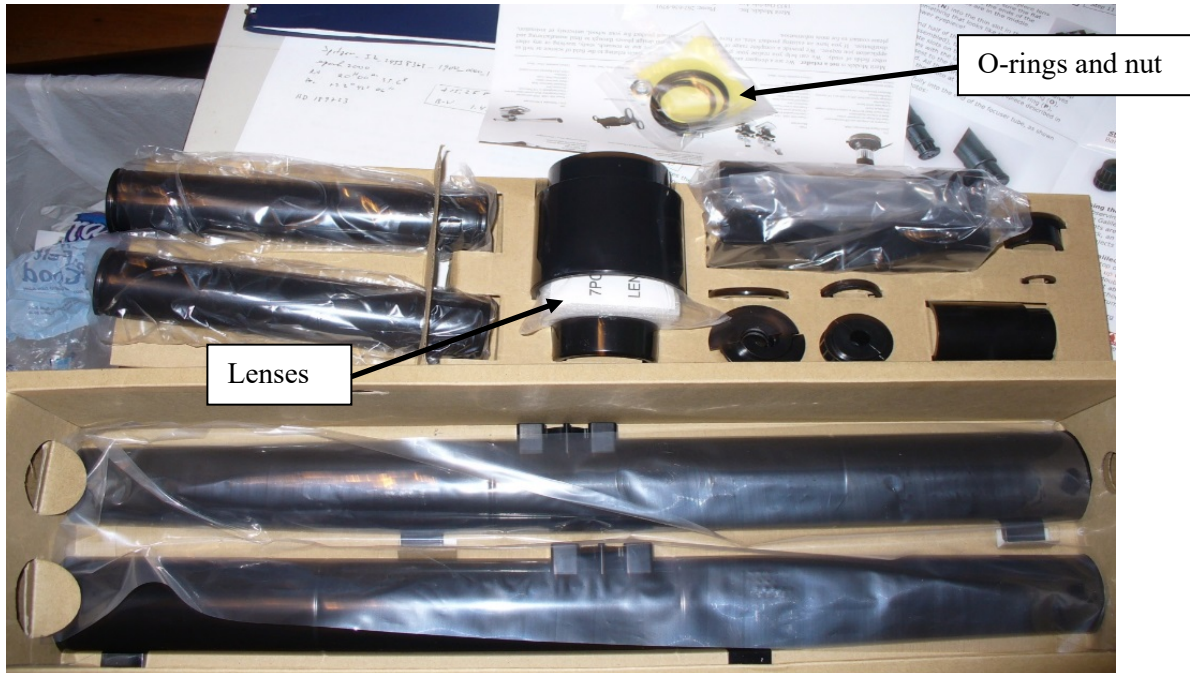
$$m = \frac{f_o}{f_e} \tag{1}$$

Galileo's original telescopes had a magnification of about a factor of 25, allowing him to observe details in the sky that had never been seen before.

Lab Procedure

Open the Galileo Telescope box and remove the top cardboard parts holder. Remove the second layer cardboard tray and set to one side.

The telescope parts should look like this:

Take careful note of where all the parts are located and how they fit into the cutouts in the cardboard. You must return the parts to their proper places for the next class to use.   You must also have your instructor check that all of the pieces are in the proper place.

Be sure to handle the lenses only by their edges, with a piece of tissue paper or cotton gloves, if available.

Parts List (in order of assembly):

A – telescope main tube halves (2)
B – V-block bases/stands (2)
C – 50-mm glass objective lens
D – ¼-20 tripod nut
E – focuser tube halves (2)
F – small main-tube clamp ring
G – small rubber O-rings (2)
H – Sun-warning sticker
I – large lens shade/dew cap
J – large rubber O-rings (2)

K – main eyepiece barrel halves (2)
L – auxiliary eyepiece barrel halves (2)
M – small main eyepiece lenses (4)
N – tiny, thin eyepiece ring/field stop
O – large main eyepiece clamp ring
P – small eyepiece clamp rings (2)
Q – tiny auxiliary eyepiece lenses (2)
R – Barlow lens tube
S – auxiliary eyepiece cap

Step 0.

Measure the diameter D of the objective lens and its focal length f in mm. You can measure the focal length by pointing the lens out the window and forming an image on a piece of paper behind the lens. Remember (from last week's lab!) that the distance to the object d_o, the distance to the image d_i, and focal length f are related by the formula

$$\frac{1}{f} = \frac{1}{d_o} + \frac{1}{d_i} \tag{2}$$

If d_o is very large (i.e., if you are looking out the window) you may use a simplified version of the formula to determine f, namely,

$$f \approx d_i \tag{3}$$

Question 3: *What is the focal length f of the objective lens?*
***Question 4:*** *What is the diameter D of the objective lens?*

Step 1.

Lay the telescope main tube halves (A) on the table, perhaps setting one of them in the V-block stands (B). Note that the objective lens (C) is actually two lenses cemented together. One lens is thinner, and the other thicker. Insert the objective lens into the groove at the end of one telescope main tube half so that the thinner lens points forward, out of the telescope as shown at right.

Step 2.

Insert the ¼"-20 tripod nut (D) into the slot in the telescope main tube half. To seat the nut securely, make sure it is oriented with one of its "points" (not one of its flat sides) facing up, as shown in the photo at right.
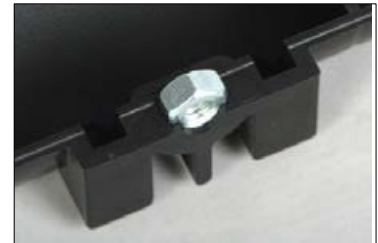
Step 3.

Lay the two focuser tube halves (E) on the table, oriented with their interiors facing up. Note that one end of each tube is rough on the inside, and the other end is smooth. On one tube half, the smooth end has two U-shaped cutouts; they're at the bottom left in the photo at left. Orient the tube halves so that the two smooth ends match and the two rough ends match, as shown in the photo at right.

Steps 4 & 5.

Join the two focuser tube halves and hold them together. Slide the small main-tube clamp ring (F) onto the focuser tube, with the wider end of the ring facing away from the end of the tube with the two U-shaped cutouts. Secure the two ends of the focuser tube with the two small rubber O-rings (G), which fit in the groves around each end of the tube. The O-rings look too small to fit, but they will stretch.

Step 6.

Lay the completed focuser tube assembly into the back end of the main tube. As shown below, make sure that the end of the focuser tube with the two U-shaped cutouts is protruding out the back of the telescope main tube, along with the main-tube clamp ring. The other end of the focuser tube should lie between the two baffles at the back of the main telescope tube as shown in the photo below.

Baffles
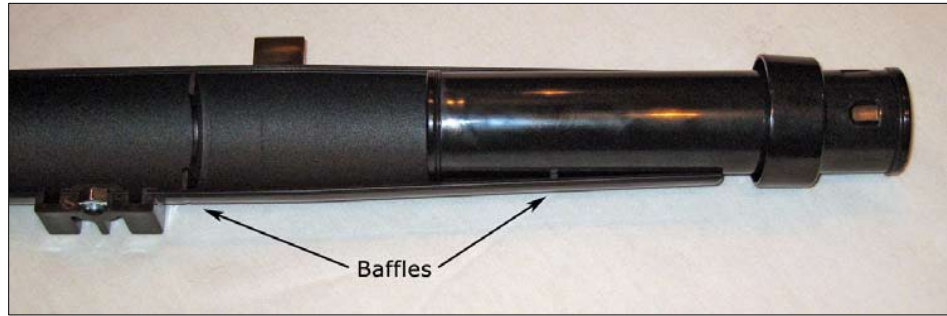
Step 7.

If not already done, peel the backing off the Sun-warning sticker (H) and affix the sticker to the second half to the telescope main tube (A), about 25mm (1 inch) from the narrow end as shown in the photo below.

DO NOT LOOK AT THE SUN! SEVERE EYE DAMAGE WILL OCCUR
**(as well as damage to the optics of the telescope).**

Step 8.

Place the second half of the main tube over the first half (the one already on the table or in the V-block stands), as shown below. Make sure the objective lens and the ¼-20 tripod nut fit securely into their slots in the top half of the tube.

Step 9.

Secure the body of the telescope together by sliding the main tube clamp ring (F) onto the back. Place the two large O-rings (J) around the telescope main tube in the channels provided for this purpose (see arrows in photo below). These will hold your telescope together more securely. The O-rings <u>will</u> stretch enough to fit over the tube. If you are unsure, please ask your instructor.
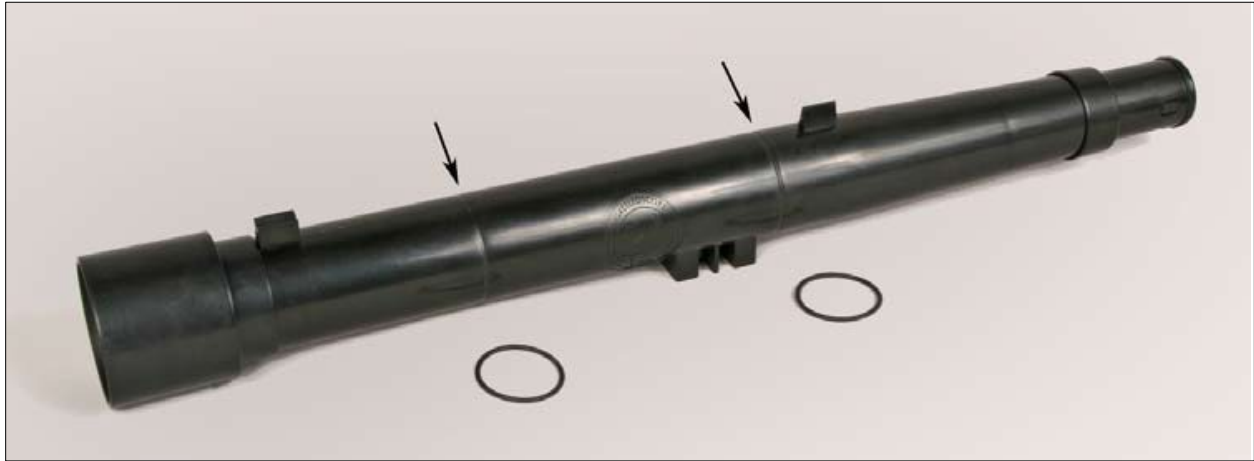
Step 10.

As seen above, attach the large lens shade/dew cap (I) onto the front.

There are two eyepiece barrels. The wider eyepiece (K), with the larger central opening is the main (or "modern") eyepiece. We will be using both for this lab, to determine how the different eyepiece designs affect the magnification and image quality of the telescope.

The "Modern Eyepiece"

We'll start out using a modern eyepiece consisting of a combination of several lenses. A complicated "compound" eyepiece design such as this yields excellent images but was not practical for Galileo because the lenses available to him contained so many flaws that, the more lenses used, the worse the image.
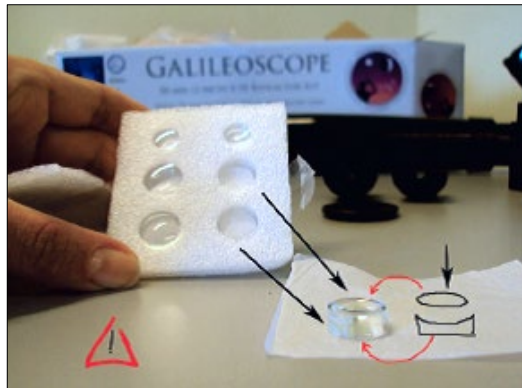
Step 11.

Open the lens packet, shown at right, and note that there are 6 lenses included. For the first experiment we will use the four larger lenses in the bottom two rows.
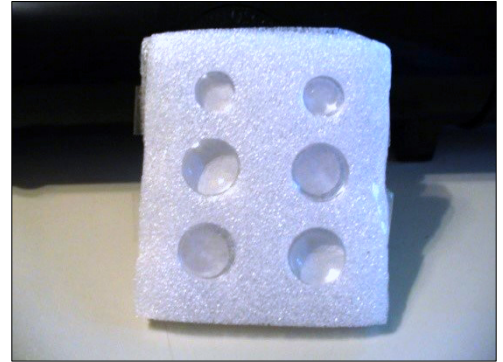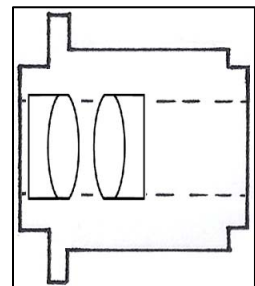

Step 12.

The four modern eyepiece lenses (M) are about 14 mm (a little over a half inch) in diameter. Handle the lenses with the supplied tissue paper, cotton gloves or touch only the edges to avoid fingerprints. Two lenses are flat on one side and concave – curved inward – on the other. The other two lenses are convex – curved outward – on both sides. Take one of each type of lens and place them together as show in the drawing. Repeat with the other modern eyepiece lenses.

Note: While wearing the cotton gloves,
you can feel for the curvature of the lenses with your fingers to pick the correct lenses.

Step 13.

Take one half of the modern eyepiece barrel (K). Insert the two eyepiece lens pairs (from step 12) into the appropriately-sized slots of the barrel. **Be sure the flat sides of the lens pairs point away from each other** (see picture on right).

Step 14.

Insert the tiny, thin eyepiece ring/field stop (N) into the thin slot in the modern eyepiece barrel half. You should now have something that looks like the photo at right. You are almost done with the modern eyepiece.

ring/field stop (N)

Step 15.

Join the second half of the modern eyepiece barrel (K) with the first half (the one you just assembled), taking care that the lenses and field-stop ring fit into the appropriate slots on the second half as you bring the halves together. Secure the halves with the large modern eyepiece clamp ring (O), which goes on the other end. The completed modern eyepiece is shown in the photo at right.

Step 16.

Measure the focal length of this compound eyepiece by forming an image of some distant object on a piece of white paper, as you did for the objective in Step 0.

    Question 5: *What is the focal length f_e of the compound Modern Eyepiece?*

Step 17.

Insert the eyepiece fully into the end of the focuser tube, as shown in the following sequence of photos:

Step 18.

Observe a scene through the lab window. Focus the telescope by sliding the focusing tube (E) back and forth to achieve a sharp image. (The back of the eyepiece should end up about 2-1/2 inches behind the end of the telescope barrel.) Note the properties of the images you see. Measure the magnification of the telescope with the modern eyepiece. You may do this by eye or with the camera on your cell phone, with and without the telescope.

Question 6. *Describe the quality and the orientation of the images you observe through the Modern Eyepiece. Was it easy to find your target? Was the image sharp? (Remember to focus the telescope!) Are the images upright or inverted?*

Question 7. *What magnification did you observe with the Modern Eyepiece? Explain how you arrived at this value.*

Question 8. *What is the "theoretical" magnification of the telescope using the modern eyepiece? (I.e., the value based on **Equation 1** and your measurements of f_o and f_e). Does it agree with your observed value from Question 7? Which one do you think is more accurate? (Show all steps in your calculation.)*

The "Galilean Eyepiece"

Now we'll try a simple single-element eyepiece, like one that Galileo used. (Except that this one is made of plastic!) It consists of a single small concave lens which yields upright images as shown in Figure 1.

Step 19.

The Galilean eyepiece uses one of the smaller lenses (about 10 mm in diameter) located in the top row of the lens packet (shown at right). The lens we will use is the one that is concave (inward curve) on both sides. You can feel for concavity by touching both surfaces <u>with the lens paper or cotton gloves covering your fingers.</u>
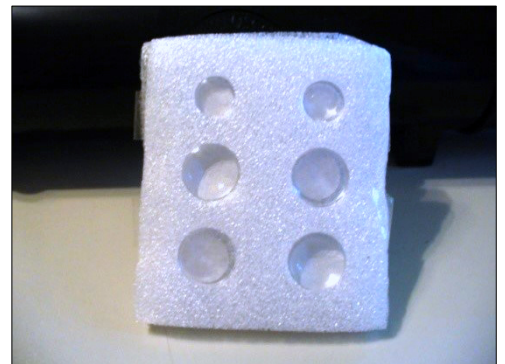
Step 20.

To make the simple "Galilean Eyepiece," first lay half of the auxiliary eyepiece barrel (L) on the bench and gently place the 10 mm, double concave lens in the grove at the center of the barrel. Make sure it is standing straight up, perpendicular to the barrel. This eyepiece can accommodate 2 lenses in this location, so be sure it sits squarely in the grove. Then join the two halves of the auxillary eyepiece and place the auxiliary eyepiece cap (S) over the barrel, as shown below, and insert into the telescope. With this eyepiece, your telescope will function like that shown in **Figure 1**.

Step 21.

Once again observe a scene through the lab window. Move the focusing tube (E) back and forth until you get the best image. (The back if the eyepiece will sit about 1-1/2 inches behind the telescope.) Don't expect a high quality image. When Galileo made his telescope, the art of lens making was not very well perfected and lenses included many air bubbles and usually had a green tint from impurities. Note the properties of the image you are observing. Estimate the magnification of the telescope using the Galilean Eyepiece. You may do this by eye or with the camera on your cell phone, with and without the telescope. Galileo's original telescope had about 20x magnification.

> ***Question 9.*** *Describe the quality and the orientation of the image. Is it easier or harder to use this eyepiece compared to the Modern Eyepiece? How does the image quality compare with that seen through the Modern Eyepiece? Which eyepiece would you rather use for your own observations? Why?*

> ***Question 10:*** *What magnification (m) do you observe with this configuration? Explain how you arrived at this value.*

> ***Question 11:*** *What is the focal length of the concave Galilean Eyepiece? Since you cannot directly measure this quantity for a concave lens, you'll have to estimate the focal length indirectly. I.e., use* ***Equation 1*** *– you already know two of the three quantities in that equation. (Show all steps in your calculation.)*

<u>Cleanup</u>

Please carefully disassemble the telescope and replace all parts in their proper locations.
<u>Do not close the box until your instructor has checked the inventory.</u>

Lab Report Cover Page

Spring 2022
Mendel Science Experience 2151
Astronomy Laboratory - Stars

Lab E
<u>Building a Galileoscope</u>

Name: _____

Lab Section: _____

Date: _____

Lab F
Classification of Stellar Spectra

(Adapted from *Contemporary Laboratory Experiences in Astronomy*)

PURPOSE:

To become familiar with the appearance of the spectra of main sequence stars.   To learn how to classify the spectra of main sequence stars by comparing a spectrum with an atlas of spectra of selected standard stars.

EQUIPMENT:
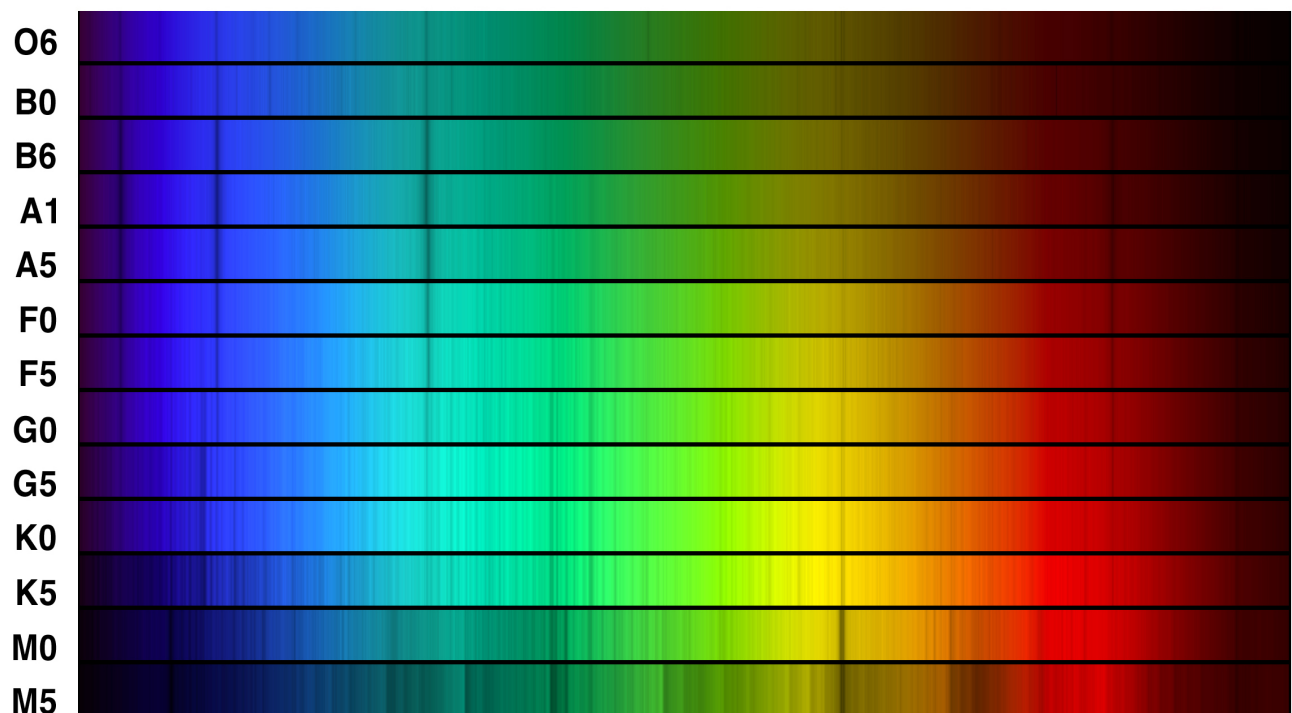
CLEA computer program ***Stellar Spectra***

Figure 1: The OBAFGKM Spectral Sequence
(Photo credit: www.noao.edu)

Spectral Classification

The German physicist Joseph von Fraunhofer was, in the early 1800's, the first to notice that the spectrum of the Sun was "missing" light at very specific wavelengths. These features, now known as the *Fraunhofer absorption lines,* were then of unknown origin and were simply labeled with letters of the alphabet – as shown in Figure 2 below.
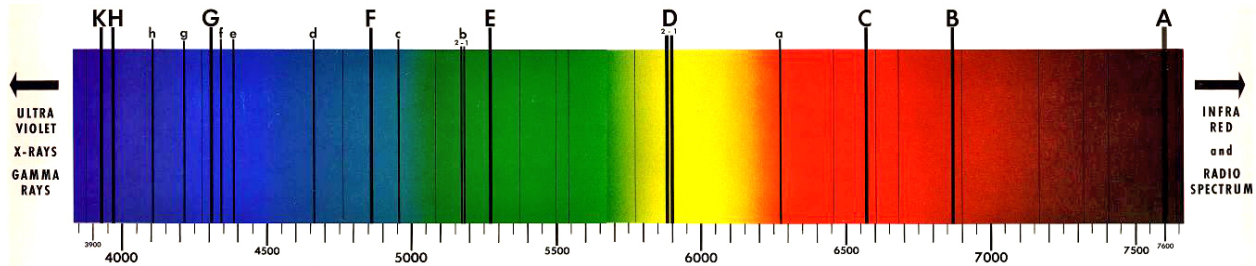

Figure 2: *The Fraunhofer lines in the Sun's spectrum*

It was not until late in the 19th century that astronomers were able to routinely obtain and examine the spectra of other stars in large numbers. The astronomers Fr. Angelo Secchi and E.C. Pickering were among the first to notice that the patterns of absorption lines varied among the stars, but that distinct groups or "types" could be identified.   This gave birth to the process of *spectral classification,* in which stars are grouped together based on the prominence of certain specific absorption lines.   Building upon this early work, astronomers at the Harvard Observatory refined these "spectral types" and named them with letters *A, B, C,* etc.   They also embarked on a massive project to classify spectra – carried out by the trio of astronomers, Williamina Fleming, Annie Jump Cannon, and Antonia Maury – which ultimately lead to the classification of over a quarter-million spectra.

The early classification systems were based strictly on the appearance of the spectra, with no understanding of the physical reasons for the differences among stars. This changed in the 1930's and 1940's when it came to be understood that the strength of a specific absorption line depends on the abundance of the type of atom which absorbs at that wavelength, which – in turn – depends on the chemical composition of the star and the temperature and density of its outer layers, where the absorption takes place. Since it turns out that most stars have very similar chemical composition, ***the surface temperature of a star is the main determinant of the appearance of a stellar spectrum.***

The spectral classification system in widespread use today is a refinement of the work by the Harvard astronomers and is called the ***MK System***.   It was introduced in the 1940's and 1950's by W.W. Morgan and P.C. Keenan of Yerkes Observatory. This system has seven main *temperature classes*, ***OBAFGKM***, which are ordered by decreasing temperature. The ***O*** stars are the hottest stars (with temperatures around 40,000 K) and the ***M*** stars are the coolest (with temperatures around 3,000 K). As a further refinement, the major temperature classes are subdivided into subtypes, e.g., B0, B1, B2, … B9.   These are also ordered by temperature so that, for example, a B0 star is hotter than a B5 star, and a B9 star in the sequence is followed by a cooler A0 star. Sample spectra are shown in **Figure 1** on the front page of the lab. Notice the systematic change in the appearance of the absorption lines moving down the spectral sequence from the *O* types to the *M* types. The ***MK System*** adds a second dimension to the classification scheme to take into account the fact that stars with the same surface temperature can have different sizes. A difference in size between two stars with identical surface temperatures results in a difference in their intrinsic luminosities and can introduce subtle differences in their spectra – which can be discerned by a skilled classifier. A Roman numeral is added to the end of an *MK* spectral type to indicate a star's *luminosity class:* an ***I*** indicates a supergiant star, a ***III*** indicates a giant star, and a ***V*** indicates a smaller Main Sequence star.   (There are

also intermediate types *II* and *IV*.) Our Sun is an ordinary Main Sequence star and has a classification in the *MK System* of *G2 V*.

The spectral type of a star is so fundamental that an astronomer beginning the study of any star will first try to find out its spectral type. Detailed analyses performed on many stars have shown that precise determinations of a star's surface composition, temperature and density can be made from study of high-quality spectra, combined with our current understanding of atomic physics. These fundamental stellar properties have been seen to correlate with stellar spectral type. Since stars which have similar spectra have similar physical properties, the properties of an unknown star can be quickly inferred from past detailed studies of other stars of similar type. A good spectral type thus provides a shortcut to knowledge of a star's fundamental properties.

Lab Procedure

In this lab you will become an expert stellar classifier. Specifically, you will examine the spectra of 20 unknown stars, determine the spectral type of each star, and record your results along with the reason for making each classification. You will compare your unknown spectra with a representative atlas of standard spectra and, by looking at the relative strengths of characteristic absorption lines, you will be able to estimate the spectral type of the unknown stars to about a tenth of a spectral class, even if they lie between the spectral types of the stars in the atlas. To make life a little easier, we will at first confine ourselves to the classification of Main Sequence stars (luminosity class *V*).   Near the end of the lab, you will examine the effects of luminosity on the appearance of the stellar spectra.

Part I - ENTERING STUDENT ACCOUNTING INFORMATION

Launch the ***Stellar Spectra*** program in the "MSE Astrolab" folder. Click ***File, Log In*** to activate the Student Accounting screen. Enter your name (first and last). Do not use punctuation marks. When all the information has been entered to your satisfaction, click ***OK*** to continue, and click ***Yes*** when you are asked if you are finished logging in. The opening screen of the Spectral Classification lab will then appear.

Part II - SPECTRAL CLASSIFICATION OF MAIN-SEQUENCE STARS

Step 1.   Select ***File, Run, Classify Spectra***. You are now in the classification tool. The screen that you see shows three (currently empty) panels, one above another with some control buttons at the right and a menu bar at the top. The center panel will be used to display the spectrum of an unknown star, and the top and bottom panels will show you spectra of standard stars that can be compared with the unknown. Let us now run through the features of the classification tool by classifying the first of the 20 unknown spectral provided for practice.

Step 2.   Call up the spectrum of a practice "unknown" star by clicking on the ***File, Unknown Spectrum, Program List.*** A window will appear displaying a list of practice stars by name. Highlight the star HD 124320, then click ***OK***. You will see the spectrum of HD 124320 displayed in the center panel of the classification screen, as shown in **Figure 3**.

Until recently, spectra were displayed as photographs of the stellar spectrum, as in **Figures 1** and **2**, but modern spectrographs yield direct measurements of the intensity of light as a function of wavelength and it is now common practice to display a spectrum as a plot of intensity (or brightness) of light versus wavelength, as in **Figure 3**. You can compare the two different display methods by choosing ***File, Preference, Display, Comb. (Photo & Trace).*** The classification tool will now display the HD 124320 spectrum in both the photograph format (without the color) and the intensity format. These convey the same
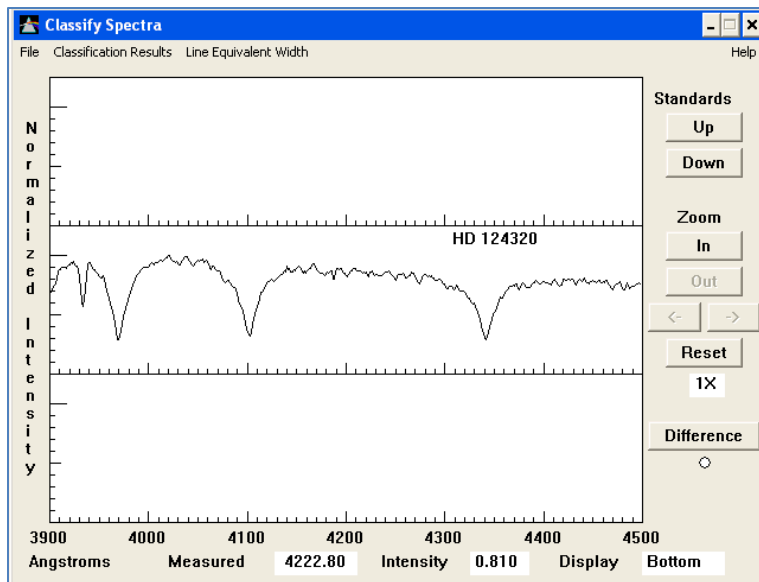
Figure 3: *The spectrum of HD 124320*

information but in very different ways. In both cases, wavelength is displayed in units of Ångstroms (Å), where 1 Å = 10^{-10} meter. In other popular wavelength units, this would correspond to 390 to 450 nm (nanometers) or 0.39 to 0.45 μm (microns). This is only a portion of the total optical spectral range (~3500–7000 Å) but contains sufficient spectral features to allow precise and accurate classifications. The intensity can range from 0 (no light) to 1 (maximum light). The highest points in the spectrum, called the *continuum*, are the overall light from the incandescent surface of the star, while the dips and dark features are the *absorption lines* produced by atoms, ions, or molecules further out in the atmosphere of the star. You can measure both the wavelength and the intensity of any point in the spectrum by pointing the cursor at it and clicking (and holding) the left mouse button. (Try it.) The cursor changes from an arrow to a cross, making it easier to center on the point desired. The "strength" of an absorption line is often gauged by its depth, i.e., the deeper the line, the stronger.

Decide whether you want to do your classifications using the intensity format (which is recommended) or the photograph format. If the former, select ***File, Preference, Display, Intensity Trace.*** If the latter, select ***File, Preference, Display, "Grayscale" Photo.***

Step 3. Now you want to find the spectral type of HD 124320 by comparing its spectrum with spectra of known type. Call up the comparison star atlas by clicking the ***File, Atlas of Standard* Spectra** option. A window will open up on which you will see numerous choices, but the atlas you want is the one on top of the list, ***Main Sequence***. Select it and click on ***OK*** to load the atlas.

Step 4. The 13 spectra in the Atlas will come up in a separate window – see **Figure 4** – but only 4 can be seen at one time. You can look at the entire set by dragging up and down on the scrollbar at the right of the Atlas window. Do this, and note that a sequence of representative types, spanning the range from the hottest to the coolest, are shown. The wavelength range, though not indicated, is the same as in **Figure 3**. As you scroll through the standard stars, notice how the shapes and strengths of the absorption features vary systematically as you move from hotter types to cooler types, as shown in the set of photographic spectra on page 1 of the lab. You should also notice that the shape and slope of the background spectrum upon which the absorption lines are superimposed (i.e., the *continuum*) also varies systematically with spectral type and, thus, stellar temperature. This is a consequence of the *blackbody radiation laws*, specifically *Wein's Law*, and will be discussed in a later lab.
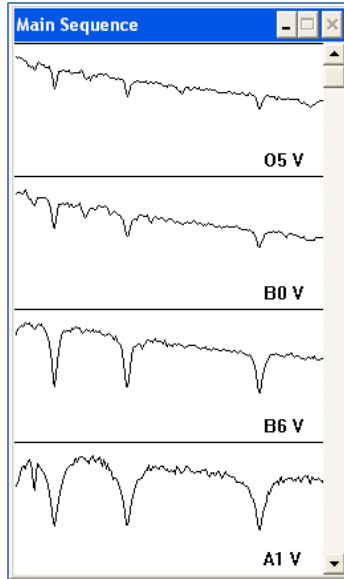
Figure 4: *The Main Sequence spectral atlas*

Step 5.   Now let's use the comparison spectra to classify HD 124320. Minimize the Atlas (if you want to restore it, you can always double click on its icon) and look at the panels in the classification tool. You will see that two of the comparison star spectra have already been placed in the two panels above and below the spectrum of your unknown star. You should see the spectrum of an *O5* star is in the top panel, and the spectrum of the next star in the atlas, a *B0*, in the bottom panel.

If neither of these looks quite like a match to your unknown star, you can move through the Atlas by clicking on the button labeled ***Down*** to the upper right of the spectrum display. Continue this until you get a pretty good match. You should find that the best match is with spectral types that have very strong hydrogen lines (more about how to identify these will come later), and not many other features. The stars with the strongest hydrogen lines are around spectral type *A0*.

Because not all spectral types are represented in the Atlas, and because you want to get the classification precise to the nearest 1/10 of a spectral type (*i.e., G2*, not just *G*), you may have to do some interpolation. Look at the relative strengths of the absorption lines to do this. For your unknown star, for instance, you should note that the strong lines at 3970, 4101, and 4340 Å are similar in strength to both the *A1 V* and *A5 V* standard star spectra, but that the feature at 3934 Å is intermediate in strength between those two.   Thus, the spectral type of HD 124320 likely lies between those of the two standard stars and is probably *A2*, *A3*, or *A4*. If you think the strength of the 3934 Å line is exactly in the middle, choose *A3*, if you think it is closer to one standard or the other, then choose *A2* or *A4*.

Step 6.   If you want to do this in a more quantitative fashion, click on the button labeled ***Difference*** to the right of the spectrum display. The bottom panel graph will change, showing the digital difference between the intensity of the comparison spectrum at the top and the unknown spectrum in the center, with zero difference being a straight horizontal line running across the middle of the lower panel.

Look at the dips and valleys on this bottom panel and think about them for a minute. If an absorption line in the comparison star is shallower than the line at the same wavelength in the unknown star, then the intensity at those wavelengths in the comparison star will be greater than those in the unknown. So the difference between the two intensities will be greater than zero, and difference display will show an upward

bump. If the top panel is showing an *A1* spectrum, for instance, and middle panel HD 124320, you should see a small bump at 3933 Å, indicating that the absorption line in the unknown is deeper than that in the A1 standard.

By the same reasoning, if an absorption line in the comparison spectrum is deeper than one in the unknown star, then the difference display will show a downward *dip*.   Click on the Standards ***Down*** button to display an *A5* comparison spectrum.   Note that the 3933 Å difference display now shows a dip, indicating that the absorption line in the unknown is shallower than that of an *A5* standard.   So it is somewhere in between an *A1* and an *A5* in classification.

To use the difference display, page through the comparison spectra (using the ***Up*** and ***Down*** buttons) until the difference between the comparison and unknown star is as close to zero at all wavelengths as possible. To estimate intermediate spectral types, watch to see when the display changes from bumps for some lines, to dips.   Since some lines get stronger with temperature, and others get weaker, you will see some lines go from bumps to dips, and some from dips to bumps, as you change the comparison spectra.   Try to estimate whether the amount of change places the unknown halfway between those two comparison types, or if it seems closer in strength to one of the two comparison types that it lies between.

Step 7.   Record your choice for HD 124320 and your reasons in the data table provided below. **Indicate which absorption lines lead you to your answer.**   For example: *"The strength of the lines at 3970, 4102, and 4340 Å are almost exactly those of type A1 or A5, and the strength of the 3934 Å line appears to lie midway between them."*

Step 8.   You have used one or more spectral lines for making a refined classification.   But what elements produced them? For reference, you will want to identify the source of the line you are looking at.   Select ***File, Spectral Line Table***. You will see a window containing a list of spectral lines, as in **Figure 5** below. Drag this window to the side so that it doesn't obscure the Classification Tool window. Now, using the mouse, point the cursor at the center of any line in the spectrum (say the one at 4102 Å) and double-click the left mouse button. A vertical red line should appear on the screen and, if you've centered correctly, a double-dashed line on the line list to show you what the spectral line is. For instance, the line at 4102 Å is a line from neutral hydrogen (H I). See **Figure 5**. Verify this identification with your own Classification Tool.

> ***Question 1:***   *What elements are responsible for the strong lines at 3934, 3970, 4102, and 4340 Å seen on the spectrum of HD124320?   What are their precise wavelengths?*

Step 9.   You have now classified one spectrum. Call up the next unknown spectrum by selecting ***File, Unknown Spectrum, Next on List***. You do not have to reload the spectral atlas. Use the methods you have practiced above to classify the remaining 19 stars on the list. Use the data table to record your results and your reasons. Be sure to identify the specific absorption lines which lead you to your classification.

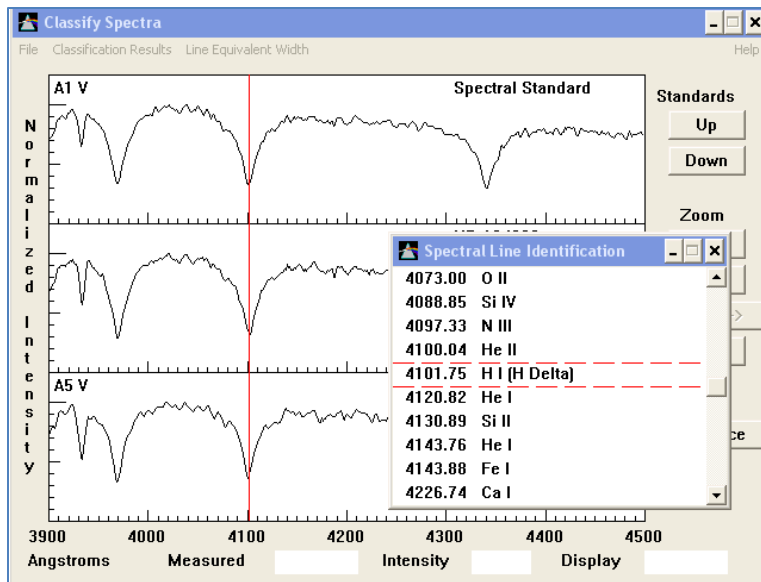***NOTE: See the Classification Hints on page D-8!***

Figure 5: *Spectral line identification*

Part III – LUMINOSITY CLASSIFICATION

Now that you have become skilled at stellar temperature classification, let's take a brief look at luminosity classification.

Step 1.   Bring up the spectrum of the practice star HD 24189. Also, click ***File, Atlas of Standard Spectra*** and choose ***Luminosity at F5***. This will load a set of standard stars that all have temperature types of *F5*, but different *MK System* luminosity classes (i.e., roman numerals *I, II, III,* etc. (See the Introduction!) Assuming that HD 24189 is an *F5* star, try to determine its luminosity class.

> ***Question 2:***   *What is the luminosity class for HD 24189? How certain are you? (Note: it's all right to be uncertain!) What features of the spectrum distinguish the various classes?*

Step 2.   Repeat the above process for the first practice star HD 124320 and the ***Luminosity at A3*** standard stars, assuming that *A3* is the appropriate temperature type.

> ***Question 3:***   *What is the luminosity class for HD 124320? How certain are you? What features of the spectrum distinguish the various classes?*

> ***Question 4:***   *Overall, do you think luminosity classification is easier or harder than temperature classification? Or does it depend on the specific spectral class being observed?*

Some Classification Hints

- One quick way to go through the spectral atlas, rather than using the *Up* and *Down* buttons is to open up the Atlas window and double click on a graph panel of the atlas representing the spectrum that you want to insert in the upper comparison panel of the Classification window. The atlas panel selected will be tinted blue to indicate that it is the one selected. You can then minimize the atlas again to see the Classification window more clearly.

- You can get close-up views of the spectra by clicking on the *Zoom In* button to the right of the Classification window. In zoom mode, the right and left arrows under the *Zoom* buttons can be used to pan along the spectrum to see wavelengths that are off the edges of the range of view. The *Reset* button returns to full view of the spectrum.

- When the Spectral Line List is visible, you can find a particular spectral line on a spectrum by pointing the cursor at an entry on the list and double clicking the *left* mouse button. A vertical red line will appear on the spectrum display at the wavelength of the line.

- When the Spectral Line List is visible, pointing the cursor at an entry and double-clicking the *right* mouse button will bring up a window with further information on the spectral line in question. In many cases several ions produce spectral lines at about the same wavelength, so it will not be immediately clear what ion is producing a particular absorption line. For instance, both Ca II and H I produce lines at a wavelength of about 3970 Å. But H I lines are strongest in *A* type stars, while Ca II lines are strongest in the *G* and *K* type stars. The notes provided in the spectral line information screens can thus be used to decide what ions are producing what absorption lines if you have a rough idea of the spectral type you are looking at.

Data Table: Spectral Classification

Star	Sp. Type	Reasons
HD 124320	A3	H I 3970, 4102, and 4340 Å lines similar in strength to A1 and A5 types. Ca II 3934 Å line strength midway between those of A1 and A5 types.
HD 37767		
HD 35619		
HD 23733		
O1015		
HD 24189		
HD 107399		
HD 240344		
HD 17647		
BD +63 137		
HD 66171		
HZ 948		
HD 35215		
Feige 40		
Feige 41		
HD 6111		
HD 23863		
HD 221741		
HD 242936		
HD 5351		

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Spring 2022

Lab F
<u>Classification of Stellar Spectra</u>

Name: _____

Lab Section: _____

Date: _____

Lab G
Effects of the Atmosphere on Astronomical Observations
(Adapted from *"The Hands-On Universe", Volume I*)

PURPOSE:

To learn how the Earth's atmosphere affects our ability to measure the brightness of astronomical objects.

EQUIPMENT:

TS-24 Telescope Simulator program and calculator or Excel spreadsheet.

How the Earth's Atmosphere Affects Star Brightness

One of the most fundamental measurements that astronomers make is to determine the brightness of stars. This process is called *photometry*. The brightness of a star as seen from Earth is related to the star's intrinsic energy output (its *luminosity*) and its distance from us – both of which are important things that we want to know. Measuring brightness is one of the first steps towards determining the intrinsic properties of stars (or other astronomical objects).

Unfortunately, most astronomical observations are done from the Earth's surface, which means that a stars' light must pass through a blanket of air before reaching our eyes or a telescope. As a result, a star looks fainter and less steady than if it were observed from above the atmosphere. This effect is illustrated in **Figure 1**. Astronomers describe this phenomenon as the combination of two effects: ***extinction*** refers to the overall reduction in the mean brightness of a star (as compared to the no-atmosphere case) and ***scintillation*** refers to the variation of the instantaneous brightness of the star about its mean value. Scintillation is commonly referred to as the *twinkling* of the stars. Note that the example in **Figure 1** shows the star to be an intrinsically steady light source. In reality, many stars are variable in their light output – but we're not going to consider that complication here!

In general, extinction and scintillation arise because photons of light can be scattered (i.e., bounced) off molecules in the atmosphere. Turbulent masses of air, called cells, continually move across the light path between an observer and a star. When a dense cell moves in front of the star, the light will scatter more and the star will dim, much like car lights in fog. When a less dense cell moves in front of the star, there will be less scattering and the star will appear to brighten briefly. There is always some scattering. The variability in the amount, which has a timescale of much less than a second, is what causes scintillation.
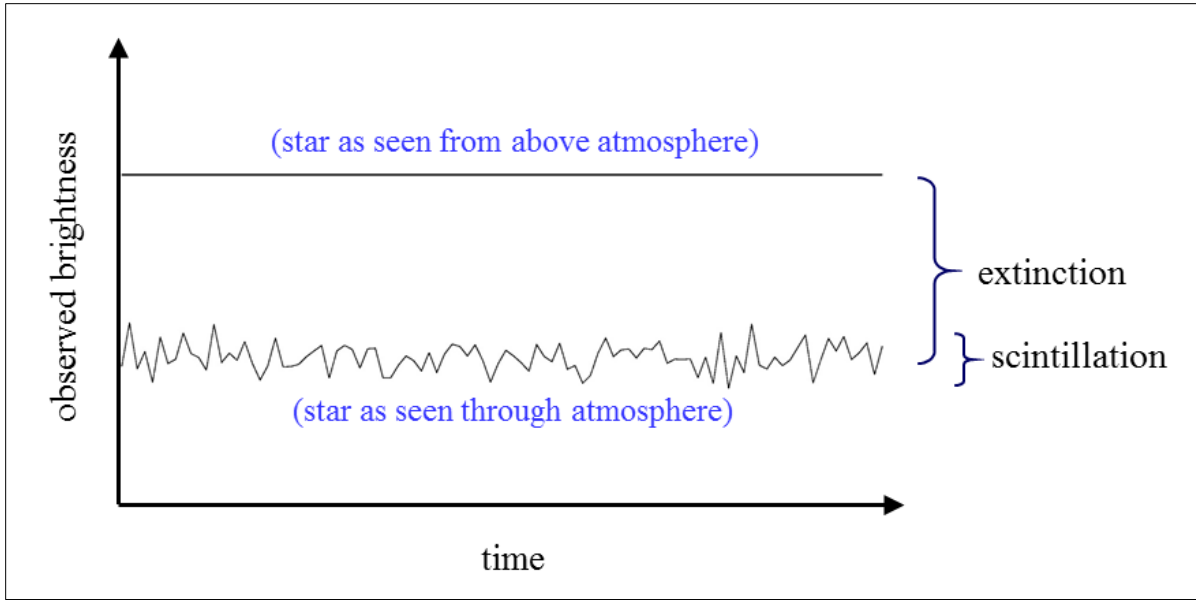
Figure 1*: The effects of the atmosphere on starlight.*

The general rule-of-thumb for astronomical photometry is that the more air you have to look through, the worse will be the effects of extinction and scintillation. We characterize the "amount" of atmosphere we look through with the quantity *airmass*, which is related to the length of the path through the atmosphere that a photon of light must travel to reach our eye or telescope. This is illustrated in **Figure 2**. It is clear that airmass is much larger for objects close to the horizon than for those high overhead, near the *zenith*. That's why the Sun looks dimmer at sunrise or sunset than at noon: the airmass is much larger and, as a result, the dimming effects of extinction are greater. The formula for computing airmass is given by:

$$airmass = \frac{1}{cos(90 - h)} \qquad (1)$$

where *h* is the *altitude* of the object, i.e., its angle above the horizon, measured in degrees (°).

The airmass is a minimum for a target seen directly overhead (at the zenith; altitude = 90°) and increases rapidly for objects close to the horizon. Note that the formula above is an approximation; it yields an infinite airmass for an object on the horizon because it assumes a flat atmosphere. Since the Earth is curved and finite in size, true airmasses never get that large. In general, because of the increasing severity of extinction and scintillation, an astronomer would be reluctant to perform photometry on an object with an airmass larger than 2.
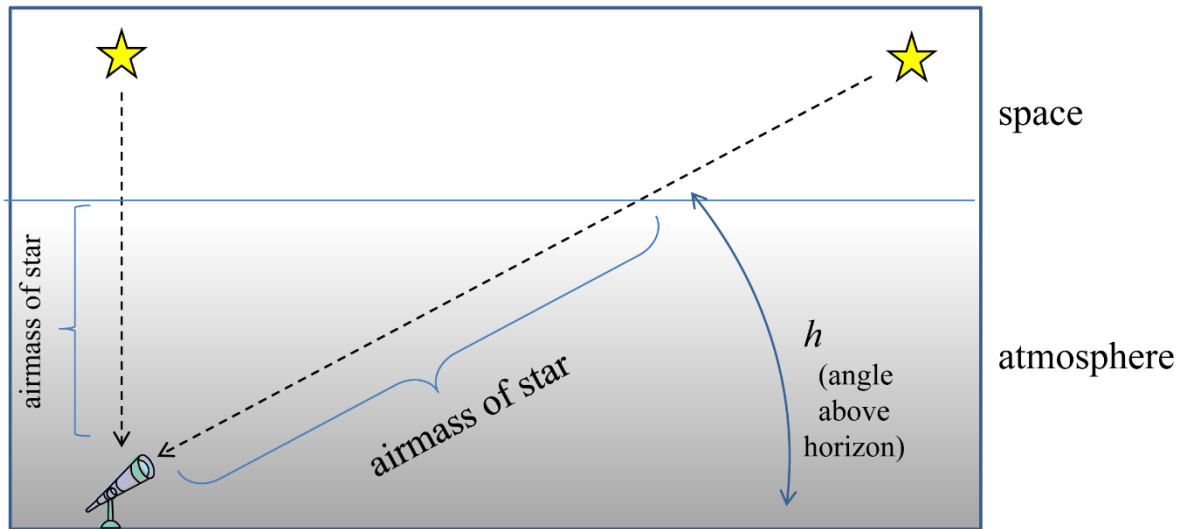
Figure 2: *A star's "airmass" depends on its angle above the horizon.*

Fill in the following table, computing the airmasses for the altitudes listed:

altitude	airmass
90°	
75°	
60°	
45°	
30°	
15°	
1°	

NOTE: if you are using Excel to calculate the airmasses, you must convert the angles into units of radians. For instance, if the angle θ is given in degrees, then the cosine of that angle is expressed in Excel as COS(RADIANS(θ)).

Question 1: Some measurement uncertainties are called "random" errors, because they cause random variations around some mean value. Other uncertainties are called "systematic" errors because they always bias a measurement in a specific direction. From the discussion above, does extinction produce systematic errors or random errors? Explain your answer. Does scintillation produce systematic errors or random errors? Explain your answer.

Lab Procedure

The goal of astronomical photometry is to make the best possible measurement of a star's mean brightness and then correct that value for the effect of extinction. In this lab we will use a simulated *photoelectric photometer* to make such measurements and explore the effects of scintillation and atmospheric extinction on the apparent brightness of a star. A photometer is a device that is mounted to the telescope and precisely measures the brightness of stars. At the heart of high precision photometers there is a *photomultiplier tube*, or PMT. The PMT is a highly sensitive device that is able to count the total number of individual photons from a distant astronomical source that enter the photometer during a specified time interval. The brightness of an object is measured as the Count Rate, i.e., the number of photons received per second.

A. EXPLORING ATMOSPHERIC SCINTILLATION

Step 1: Set up the observatory

Start the ***TS-24 Telescope Simulator*** program in the "MSE Astrolab" folder. Click on **OBS** and choose the **Villanova** Observatory. Click on **TIME**, choose Local Time (**LST**), then select 3:00 am in the morning of February 1, 1999. Click on **OK** when you are sure that the date and time are correct. Next, we need to choose a telescope to view through. Click on **TEL** and choose the **5" Finder** telescope. This telescope will make it easier to identify our target star; then we will switch to the larger telescope with the photometer attached to make our measurements

When all this is done, you should see in the center of your screen the view through the telescope. In the upper left corner are the clocks. One displays Local Time (it should read close to 03:00:00), and another displays Universal Time (UTC), which is the current time in Greenwich, England at a longitude of 0°. On the right side of the screen are the positioning controls and the coordinates towards which the telescope is currently pointed.

Click on the button that changes the Slewing Rate to **1** arc degree-increments. This will allow the telescope to move quickly from place to place on the sky

Below the view are the eyepieces that can be used with the Finder telescope. They are classified by their focal lengths **40mm**, **30mm**, **20mm**, and **10mm**. Click on a few of them to see how it affects the field of view through the instrument. You should recall from an earlier lab that the magnification seen through a telescope depends on the focal length of the eyepiece. Clicking on the "+" brings up a cross hair to help in fine tuning the position of the telescope.

Step 2: Point to our first target star, *pi¹ Boötis*

Next, we need to point the telescope to our scintillation target star. The position of the star on the sky has been stored in the Observatory's memory. To retrieve it, click on **DT-7** to activate the Digital Tracking control. Click on the group **STARS** and then on the star *pi¹ Boötis.* The coordinates displayed for this star should be:

$$RA = 14^h\,41^m\,57^s \quad Dec = 16°\,24'\,11"\,N$$

where **RA** refers to "Right Ascension" and is a measure of the star's longitude on the sky, and **Dec** refers to "Declination", the star's latitude. If the coordinates don't match, call your instructor! If they do, then click **Move To** and the telescope will begin to move toward the desired coordinates. Alternatively, the

telescope can be moved manually using the "+" and "–" buttons below the coordinate windows.

Step 3: Fine-tune the pointing

Make sure that the **40mm** eyepiece is installed, which gives the smallest magnification and the widest field of view. When the telescope finishes slewing, there should be a star near the center of the eyepiece, near the center of the cross hair. If not, check that the telescope is indeed pointing to the correct sky location. Change to the **10mm** eyepiece. You may need to slightly adjust, i.e., fine-tune, the pointing of the telescope to bring the star to the exact center the field of view. To do this, first change the slewing rate to arc-seconds. If you lose the star altogether, simply repeat Step 2.

Step 4: Set up the photometer

Now that we've got the target star where we want it, we'll switch to the 15" telescope and its attached photometer. Switch telescopes by clicking on **TEL** and choosing the **15" Photometric** telescope. You now see the view of the sky through the photometer itself. The field of view is much smaller than with the Finder telescope, and instead of a cross hair you see a circle which represents the diaphragm of the system, i.e., the hole through which light passes onto the PMT. All of the light that passes through this diaphragm will be measured by the photometer. Click on the button that changes the diaphragm size to **28'**. Other, larger-sized diaphragms are available, and would be useful for examining larger objects like galaxies or gas clouds. For observing stars, however, we want a small diaphragm to minimize the possibility of contamination of our measurements by light from nearby stars. Adjust the position of the telescope so that pi^1 Boötis is centered in the chosen diaphragm.

Choose the visual filter by clicking **V** in the filter panel. This restricts our measurements to the wavelengths of light near the middle of the range visible to our eyes.

At this point, the screen should like this:
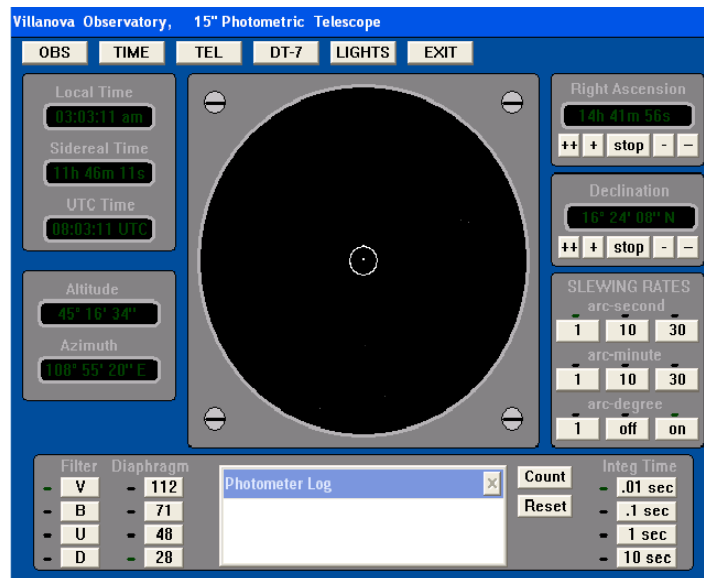
If you don't see a star centered in the diaphragm, call your instructor!

Step 5: First set of measurements

Set the integration time to **0.01 sec**. This is the length of time that the photometer will count photons of light coming from *pi[1] Boötis*. Reset the **TIME** to 3:00 am. (We want all of our measurements made at the same time so that the effects of extinction are the same.) Then measure the brightness of the star centered in the diaphragm by clicking once on the **Count** button. Numbers will appear in the Photometer Log box that list the Universal Time of the observation, Filter, Integration Time, Diaphragm Size, and Photon Count, in that order. Enter the Universal Time (UTC) and Photon Count in the first line of Table 1. The critical number here is the Photon Count, which is the number of photons that reached the photometer during the integration time. Make 9 additional measurements. For each measurement, reset the time to 3:00 am before performing the count so that all measurements are made thru the same airmass.

Step 6: Second set of measurements

Repeat step 5 above, this time with an integration time of **1.0 sec**. Record the data in Table 2.

Step 7: Third set of measurements

Repeat step 5 above, this time with an integration time of **10.0 sec**. Record the data in Table 3.

Step 8: Data analysis

For each of the 30 measurements just made, compute the photon Count Rate by dividing the Photon Counts by the integration time. Enter the result in the 4th column of Tables 1, 2, and 3.

Now compute the average Count Rate for each of the data sets. Record it on the bottom line of Tables 1, 2, and 3, and also in Table 4. **If you are using Excel**, use the *AVERAGE* function to make this computation.

Now let's look at the "scatter" in the observations. The "standard deviation" gives a quantitative description of the variability in the Count Rates. The larger the standard deviation, the larger the variability and the more uncertain your average value. **If you are using Excel,** just use the *STDEV* function to compute the standard deviation for each set of measurements, and enter the results in Tables 1, 2, and 3, and also in Table 4.

If you are not using Excel, you will have to compute the standard deviation "by hand" using the following formula:

$$Standard\ Deviation = \ \sqrt{\text{average}[(x_1 - \bar{x})^2, (x_2 - \bar{x})^2, (x_3 - \bar{x})^2, ...]} \qquad (2)$$

where the x_i are the individual count-rate measurements and $\bar{x}$ is the average count-rate.

> **Question 2**: *You should notice that the average Count Rates recorded in Table 4 for the 0.01 sec, 1.0 sec, and 10.0 sec integrations are all very similar. Why do you expect this to be the case? Describe the change in the standard deviation with integration time. For which integration time are the individual measurements most precise (i.e., show the least scatter). Considering the picture in Figure 1, explain why you expect this to be so. If you could only make one measurement, which exposure time would you choose? Why?*

As you have realized by now, the average Count Rates are simply a description of how bright a star appears to be. Astronomers use a more complicated system to describe the brightness of a star, called the "visual magnitude" system. Compute the visual magnitude corresponding to each of the average Count Rate values in Table 4 using the formula:

$$\text{Visual Magnitude} = -2.5 \times log_{10}(\text{Count Rate}) + 18.0 \qquad (3)$$

and record them in the last column of the table. This formula uses the base 10 logarithm, which is expressed as *LOG10* in Excel. Remember that magnitudes are <u>never</u> expressed in scientific notation.

B. EXPLORING ATMOSPHERIC EXTINCTION

Step 1: Set up the observatory and point to the target star

EXIT then restart the TS-24 program.   This will reset all the counters in the program and assure the most accurate answers. Choose the Villanova Observatory and set the local time to midnight on the night of February 2, 1999 – the day after the previous measurements were made. Use the 5" Finder telescope (**TEL**) and the **DT-7** to point the telescope at the bright star called *Arcturus* in the **STARS** Group. When the telescope finishes slewing make sure the telescope is pointed near *Arcturus*:

$$RA = 14^h 13^m 57^s \quad Dec = 19° 10' 15" N$$

Step 2: Set up the photometer

Switch to the 15" Photometric telescope (**TEL**) and fine-tune the position of the telescope so that the star is centered in the photometer's diaphragm. Make sure the **28'** diaphragm is chosen. Also, select the visual filter by clicking **V** in the filter panel and set the integration time to **1 sec**.

Step 3: Make measurements

 We will now observe *Arcturus* over a several hour period during the night to see how atmospheric extinction affects the light we receive. You will see that at midnight on February 2, 1999, *Arcturus* was close to the eastern horizon, as seen from Villanova. As the hours went by, Arcturus traveled higher in the sky, moving from east to west. This is due to the rotation of the Earth. Arcturus appeared close to the zenith (almost directly overhead) just before the Sun came up in the morning.

Using **Count**, measure the brightness of *Arcturus* at a local time of 0:00:00 (i.e. midnight). Record the Photon Count, and the star's altitude (angle above the horizon) in the appropriate row of **Table 5**. Round off the altitude to the nearest degree. Now make 9 more measurements, successively increasing the local time by ½-hour for each measurement. What you are doing is monitoring the brightness of *Arcturus* as it climbs higher and higher into the night sky.

When all of your observations are finished, you can close the Observatory by clicking **EXIT**

Step 4: Data analysis

Compute the airmass and visual magnitude for each of your Arcturus observations and enter the results in **Table 5**.

Using Excel, plot visual magnitude (Y-axis) vs. airmass (X-axis). The airmass axis should extend from 0

(at left) to about 4. The magnitude scale should extend from about 3 (at bottom) to about 0 (at the top). Note that, in the magnitude system, smaller values mean brighter. You should notice that the observed brightness of Arcturus drops off systematically as airmass increases and, thus, how strong the effects of extinction are! Stars are much brighter when seen high in the sky than when seen near the horizon. Your measurements (should!) show that Arcturus was almost 3 times brighter at 4:30 AM than it was at midnight.

As noted earlier, one of the goals of photometry is to <u>CORRECT</u> for the effects of extinction, not just notice that they are present. **I.e., even though our observations were made from Earth's surface, we want to estimate the brightness a star would have if it were observed from above the atmosphere.** This means extrapolating our measurements back to a value at airmass = 0. On your graph, you should see a simple linear relationship between airmass and visual magnitude. Use a ruler and extrapolate the visual magnitude back to an airmass of zero. Draw the line of extrapolation on your plot. This is your estimate of the visual magnitude of Arcturus as seen from above the atmosphere. **If you are using Excel**, use the Trendline feature of Excel to fit a linear relation to the magnitudes. Make sure to show the equation of the trendline on the graph. Use the resultant Trendline formula to compute the magnitude at airmass = 0 (this will be the intercept in your linear Trendline formula). This is the value you would measure if there were no atmosphere.

> *Question 3: What is the corrected visual magnitude for Arcturus? (I.e., the value that you would estimate for Arcturus if it could be observed from above the atmosphere, i.e., at airmass = 0). How close is this to the accepted value for Arcturus? (Go into your favorite web search engine and type something like "visual magnitude of Arcturus" to find out what the accepted value is.) Indicate what website you used for your results. You should be within about 0.5 mag of the true value.*

C. THE EFFECT OF SKY BRIGHTNESS

One final effect must be considered when trying to achieve the highest accuracy in stellar photometry – that of light <u>added</u> by the atmosphere. We have all noticed that some places have darker nights than others and, as a result, many more faint stars can be seen. This is often the result of "light pollution" in which light from terrestrial sources (such as a nearby city) is bounced around by atmospheric molecules and results in a bright background against which we are trying to see/measure faint astronomical sources. This is one of the reasons that modern observatories are located far from densely populated areas.

You have seen that the 28 mm diaphragm you used with the photometer is significantly larger than the apparent diameters of the stars being observed, thus your measurements of the brightness of pi[1] Boötis and Arcturus were actually measurements of "star + sky." In practice, astronomers try to use the smallest possible diaphragm size, but it is impossible to totally exclude light from the sky in photometric measurements.

> *Question 4: Given that the diaphragm size will always be bigger than the stellar image and that there are no places where the sky is totally dark, your measurements will <u>always</u> be contaminated. Suggest a way in which the sky brightness could have been measured and removed from your observations of pi[1] Boötis and Arcturus. (Note: Don't say "make your observations from above the atmosphere."*

Table 1: Counts for 0.01 sec Integration Time

Measurement	UTC	Photon Count	Count Rate (photons/sec)
1			
2			
3			
4			
5			
6			
7			
8			
9			
10			
Average			
Standard Deviation			

Table 2: Counts for 1.0 sec Integration Time

Measurement	UTC	Photon Count	Count Rate (photons/sec)
1			
2			
3			
4			
5			
6			
7			
8			
9			
10			
Average			
Standard Deviation			

Table 3: Counts for 10.0 sec Integration Time

Measurement	UTC	Photon Count	Count Rate (photons/sec)
1			
2			
3			
4			
5			
6			
7			
8			
9			
10			
Average			
Standard Deviation			

Table 4: Scintillation Summary Table

Integration Time (sec)	Average Count Rate (counts/sec)	Standard Deviation	Visual Magnitude
0.01			
1.0			
10.0			

Table 5: Results for Extinction Exploration

Local Time	Altitude (°)	Photon Count	Airmass	Visual Magnitude
0:00 AM				
0:30 AM				
1:00 AM				
1:30 AM				
2:00 AM				
2:30 AM				
3:00 AM				
3:30 AM				
4:00 AM				
4:30 AM				
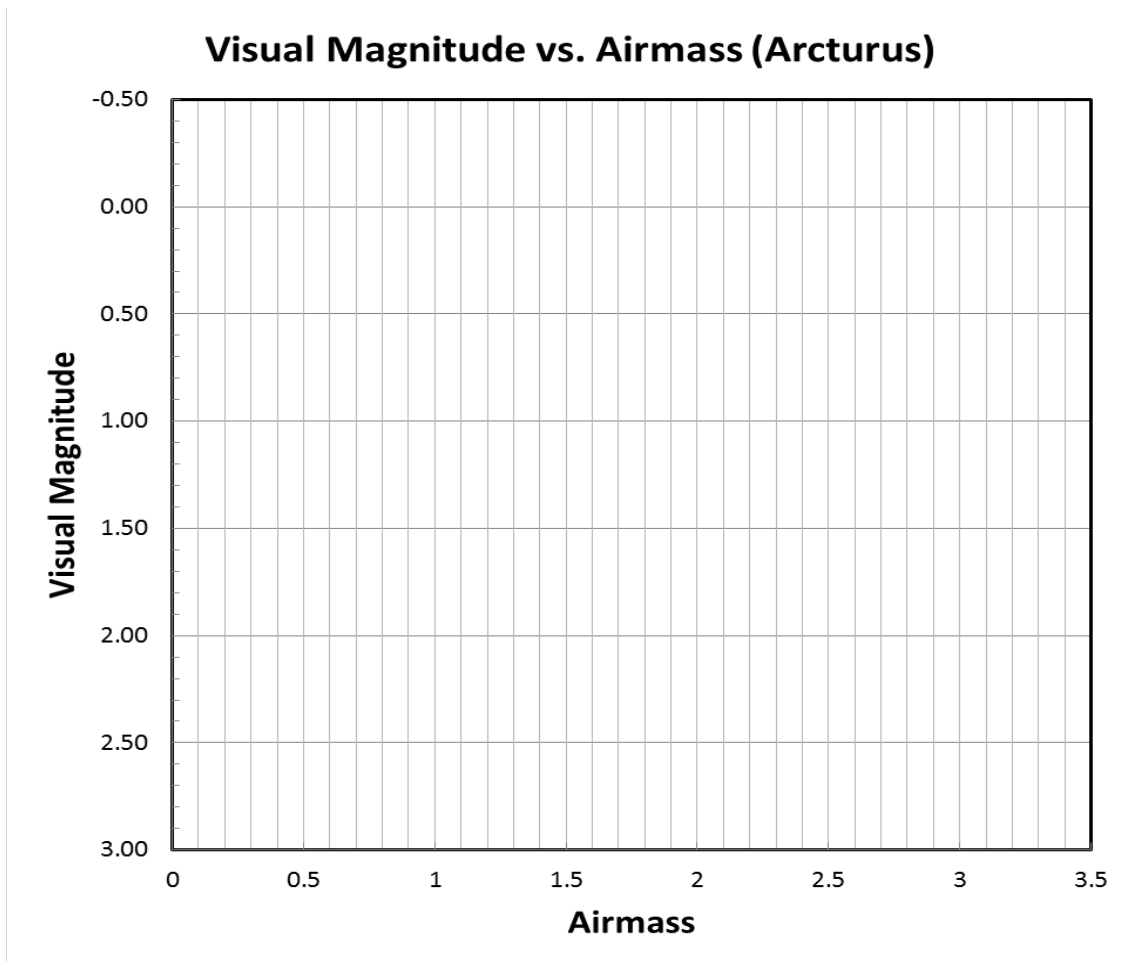
Visual Magnitude vs. Airmass (Arcturus)

Lab Report Cover Page

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Spring 2022

Lab G
Effects of the Atmosphere on Astronomical Observations

Name: _____

Lab Section: _____

Date: _____

Lab H
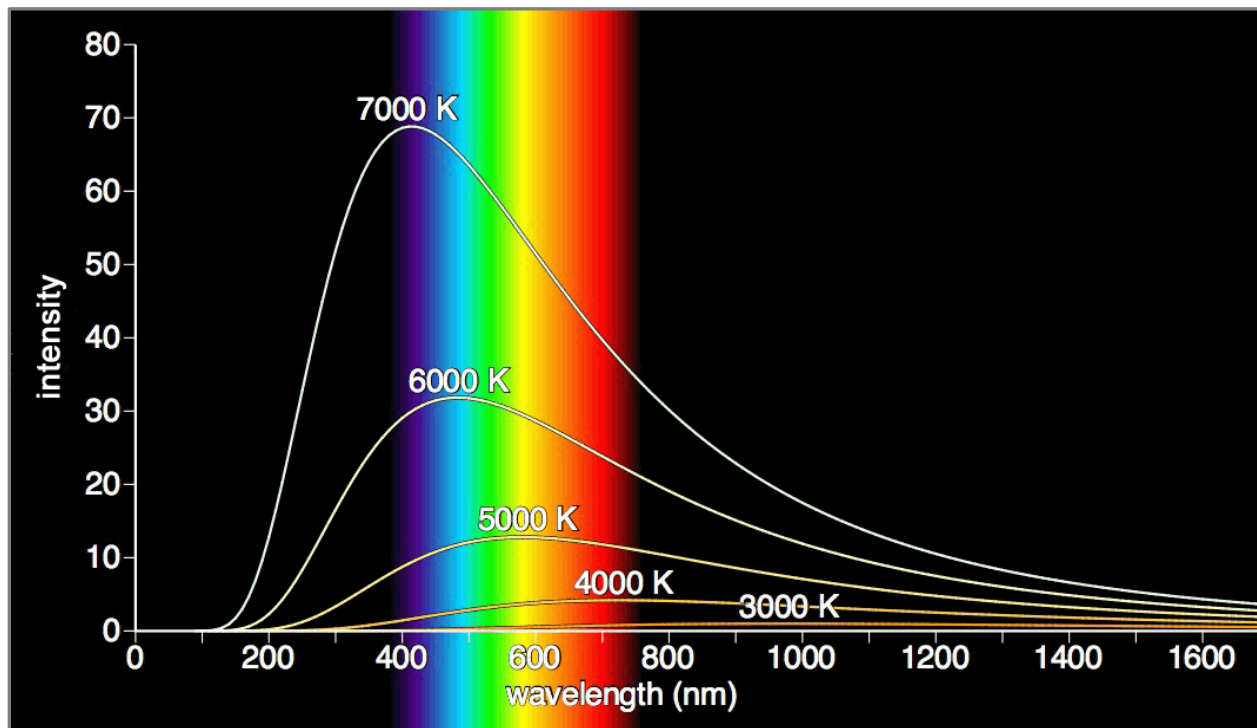## Taking a Star's Temperature
(Adapted from *The Hands-On Universe, Volume I*)

PURPOSE:

Verify that the surface temperature of a star can be estimated by measuring its brightness through a set of differently colored filters.

EQUIPMENT:

TS-24 Telescope Simulator program and calculator or Excel spreadsheet

Figure 1: *The energy distributions of hot objects*
(Picture credit: www.exo.net)

Introduction: Magnitudes, Colors & Temperature

In an earlier lab you learned that the pattern of absorption lines seen in the spectra of stars depends strongly on the temperature of their outer surface layers. Because of our understanding of the formation of spectral lines, we are able to analyze stellar spectra and determine a star's surface temperature. This knowledge is what led to the ordering of spectral types in the *OBAFGKM* classification system. The acquisition of a high-quality stellar spectrum and the detailed analysis of its properties can be a time consuming and resource-intensive process – particularly if we wish to determine the properties of many stars. Fortunately, there are other techniques available which allow astronomers to estimate stellar temperatures using more easily obtained and analyzed data. In this lab, you will examine one such technique, which utilizes stellar photometry, i.e., the measurement of the brightness of stars in specific wavelength intervals.

A hot dense object emits light at all wavelengths. In the late 1800's, the German physicist Max Planck discovered that the amount of energy radiated at wavelength λ by such an object depends only upon its temperature. He formulated a mathematical relationship which describes the brightness (or intensity) at a given wavelength as a function of temperature. This relationship is now known as ***Planck's Law***. Several examples of Planck curves are shown in **Figure 1** on the first page of the Lab, for temperatures in the range 3000 – 7000 K. The key point is that the peak in the emitted radiation shifts to shorter wavelengths as the temperature increases. This property, known as ***Wien's Law***, results in a change in the shape of the emitted spectrum within the wavelength range to which the human eye is sensitive (indicated by the rainbow). For temperatures above 7000 K, the peak shifts out of the range of human vision, although a portion of the "tail" of the energy distribution is always visible to the eye.

Stellar absorption lines are superimposed on a *continuum* spectrum which forms in slightly deeper, hotter and denser layers than the absorption lines. These layers should obey *Planck's* and *Wien's Laws* and thus the shape of the continuum spectrum should reflect the stellar temperature near the surface.

The shape of a star's continuum spectrum can be measured using "photometry." Last week you became photometry experts when you investigated the effects of Earth's atmosphere on the brightness of stars. Now you will put this expertise to work and measure the brightness of a sample of stars using a specially designed set of filters to transmit specific regions of the visible spectrum. The filters are known as the ***UBV*** filters, which stand for *Ultraviolet, Blue*, and *Visual*. The wavelength ranges which these filters transmit are shown schematically in **Figure 2**. The figure also shows schematic spectra of two stars of very different temperatures. It can be seen clearly that the ratio of the brightness of a star in any pair of filters, say, ***B*** and ***V*** or ***U*** and ***B*** will be dependent on temperature, at least over some temperature range.
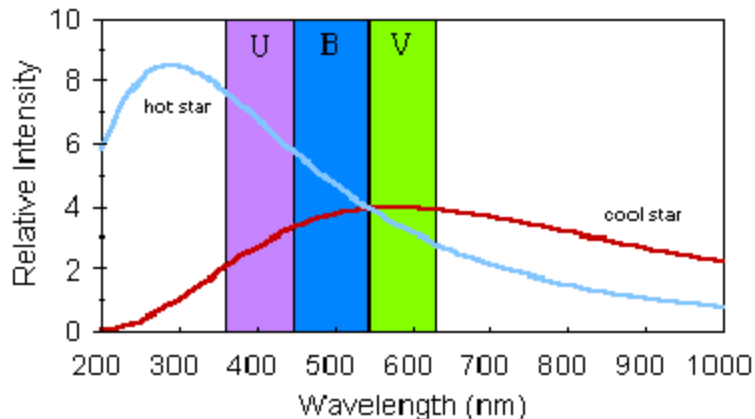

Figure 2: *Wavelength ranges of the UBV filters.*

Lab Procedure

The observations for this lab will be performed using the photoelectric photometer mounted on the 15" telescope at the *VU Special Observatory*. Why is this observatory special? Because it is located on the Moon! Why is it located on the Moon? Because – as you know from an earlier lab – Earth's atmosphere provides severe observational challenges. If we made these observations from the Earth's surface, we would have to worry about (and correct for) the effects of extinction, scintillation, and sky brightness in order to correctly measure the apparent brightness of the stars. This would add a significant amount of extra work and you would never be able to complete this lab in a reasonable time. The inconveniences of working from the Moon's surface (slight shortness of breath and intense feelings of cold) seem a small price to pay for this benefit.

You will measure the brightness of each of a set of 13 target stars using the *U*, *B* and *V* filters. These raw measurements, in units of photon counts per second, will then be converted into the standard astronomical magnitude system and you will examine the sensitivity of the "*B–V* color index" and the "*U–B* color index" to stellar temperature. The stars you will observe have been chosen to span a wide range of temperatures and to be located in a small region of the sky, so that all the observations can be accomplished in a single observing session.

A. THE OBSERVATIONS

Step 1: Go to the observatory

Open the ***TS-24 Telescope Simulator*** in the "MSE Astrolab" folder. Select the VU Special Observatory (**OBS**, **VU Special**). In the *Time* box, select Local Date/Time and set the date to November 15, 1999 at midnight local time. We've chosen this date and time so that all the stars we wish to observe will be above the horizon and high in the sky for Earth-bound observers located at Villanova.

Step 2: Prepare the telescope

Install the Finder scope (**TEL**, **Finder**) and click the *1 arc-degree* button in the *SLEWING RATES* panel, to insure that the telescope moves at its highest possible speed to each new target star.

Step 3: Observe your first star

Now perform *UBV* photometry on the first of your 13 target stars. The list of targets can be found using the DT-7 telescope control program (**DT-7**). A list of 15 stars will appear. You've already observed the first two stars on the list (pi[1] Bootis and Arcturus) when you studied the effects of the atmosphere on stellar photometry. Today's observations begin with the third star on the list, *upsilon Ori,* and end with the last star, *alpha Ori*. To move the telescope to the first star, simply click on the name *upsilon Ori* and then click the **Move To** button.

When the telescope arrives at *upsilon Ori*, install the photometer (**TEL**, **Photomultiplier**). Set the photometer parameters as follows:

<div align="center">

Integ Time: 1 sec
Diaphragm: 28 arc-seconds

</div>

Center the star in the 28 arc-second diaphragm, making sure – if necessary – to exclude any nearby stars. Then measure the brightness of *upsilon Ori* using the *U*, *B* and *V* filters by clicking on the appropriate filter

and then pressing **Count**. Record the *U*, *B* and *V* count-rates in **Table 1** or in an Excel spreadsheet supplied by your instructor. Note that, since the integration time is set to 1 second, the total number of photons counted by the photometer (last entry in each line of the Photometer Log) is the same as the count-rate in units of counts/second.

Step 4: Observe the rest of the sample

Repeat Step 3 for each of the remaining 12 stars in the DT-7 target list and record the count-rates in **Table 1**.

> **Question 1:** *The use of the VU Special Observatory just saved you a lot of time in acquiring good measurements of the brightness of your target stars. What is special about the VU Special Observatory? What benefits does it offer? What extra steps would you have had to take if you had been observing from a typical site on the Earth's surface?*

B. THE DATA ANALYSIS

Take a look at **Table 2**. The second and third columns contain basic data for your 13 target stars. The spectral types were determined by astronomers studying the optical spectra of the stars, much as you did in an earlier lab. The temperatures, which refer to the outer layers of the various stars, were determined by analyzing the strengths of the various absorption lines present in the spectra. In this section we will determine whether the overall shapes of the optical spectra are also dependent on temperature, as might be expected from *Wien's Law*.

Step 1: Compute magnitudes and colors

Compute the **V** magnitude, **(B–V)** color index, and **(U–B)** color index for each star in **Table 1** and place the results in the appropriate boxes in **Table 2** (or in your Excel spreadsheet). Use the following formulae to convert your count-rates to magnitudes and color indices:

$$\mathbf{V} = -2.5 \times \log_{10} (\text{V count-rate}) + 18.0 \tag{1}$$

$$\mathbf{(B\text{–}V)} = -2.5 \times \log_{10} \left(\frac{\text{B count-rate}}{\text{V count-rate}}\right) + 0.17 \tag{2}$$

$$\mathbf{(U\text{–}B)} = -2.5 \times \log_{10} \left(\frac{\text{U count-rate}}{\text{B count-rate}}\right) - 0.93 \tag{3}$$

The constants on the right hand side of these formulae are needed to transform your count-rate measurements into the standard system of colors and magnitudes used by astronomers. In your table, present your results to only 3 decimal places (although the count-rates are determined to a higher precision).

The **V** magnitude characterizes the brightness of a star in the night shy and depends on both the intrinsic luminosity of a star (i.e., how much energy it puts out) and its distance from us. From a good dark site, the unaided human eye can see stars as faint as about sixth magnitude (i.e., **V** ~ 6.0). Notice that nearly all of your target stars should be easily visible in the night sky. Remembering that, the fainter the star, the larger the **V** magnitude.)

> **Question 2:** *How many of your target stars would <u>not</u> be easily visible to the unaided human eye? Which ones?*

The **(B–V)** and **(U–B)** color indices are related to the shape of star's optical spectrum, and not to its brightness. More specifically, they are related to the slope of the spectrum in the region of B and V filters (for the **(B–V)** index) and in the region of the U and B filters (for the **(U–B)** index). From **Figure 2** you should expect that the indices are sensitive to the temperature of stars since the slope of the optical spectrum changes from being negative for the hottest stars (i.e., brighter at U than at V) to being positive for the coolest stars (i.e., brighter at V than at U).

Step 2: Investigate the temperature-dependence of the color indices.

Now we'll examine the sensitivity of the shape of the optical spectrum to stellar temperature. To do this, we will use the **(B–V)** and **(U–B)** color indices. Make a plot of the stellar temperature (x-axis, from column 3 of **Table 2**) versus your measurement of **(B–V)** (y-axis). Make a second plot of the temperature versus your measurement of **(U–B)**.

> ***Question 3:*** *Do your data show that the **(B–V)** color index and/or the **(U–B)** color index is sensitive to the stellar temperature? I.e., do you see a distinct relationship between the two quantities? Does this relationship make sense given Wien's Law? Refer back to **Figure 2** to remind yourself how Wien's Law affects UBV photometry. Also, take a look at **Table 1** to see which regions or the spectrum are brightest for each star.*

> ***Question 4:*** *For which stars is **(B–V)** most sensitive to temperature? The hotter ones or the cooler ones? I.e., for which stars does a small change in temperature – say a few thousand degrees – make the largest change in the observed **(B–V)**? Your observations do not include any O-type stars, which have temperatures that could range up to as high as 50,000 K. Given the data on your plot, what do you think the approximate **(B–V)** color index of an O star might be? Do you think that **(B–V)** would be a good diagnostic of the temperature of an O star? What about **(U–B)**?*

> ***Question 5:*** *Why do you think **(B–V)** and **(U–B)** are referred to as "colors" by astronomers? What do they have to do with the color of a star in the night-time sky? Neglecting atmospheric effects, do you think the color you see tells you anything about a star?*

Step 3: Compute and test the "color temperatures"

There is a standard relationship that has been used for many years by astronomers to compute the "color temperature" of a star, based on its **(B–V)** color index. This relationship is given by:

$$\text{color temperature (°K)} = \frac{7200}{\text{(B–V)} + 0.64} \tag{4}$$

and produces a temperature in units of Kelvin degrees (°K).

You will now test this relationship by comparing the color temperature for each of your stars with the spectroscopically determined temperatures.

Compute the color temperatures for all of your stars according to **Equation 4** and enter them in the last column of **Table 2**. Now make a plot of the actual stellar temperature (x-axis, from column 3 of **Table 2**) versus the color temperatures you just computed (y-axis).

Question 6: *Does the "color temperature" from Equation 4 accurately reflect the actual stellar temperature as determined from the spectral lines? If so, then the points you just plotted should all lie along a line with a slope of 1 and an intercept at 0. To check this, draw such a line on your plot. (If you used Excel to make your plot, print it now – in portrait mode – and adds this line by hand using a ruler.) Over what stellar temperature range would you feel safe in using the color temperature to estimate the actual stellar temperature? What range of spectral types does this correspond to?*

Table 1: Observational Data

Star	U count-rate (cts/sec)	B count-rate (cts/sec)	V count-rate (cts/sec)
upsilon Ori			
lambda Lep			
HD 31726			
eta Aur			
tau Ori			
beta Ori			
mu Ori			
iota Tau			
gamma Lep			
lambda Aur			
tau Aur			
rho Ori			
alpha Ori			

Table 2: Magnitudes, Colors, & Temperatures

| Star	Spectral Type	Temperature (˚K)	V	(B-V)	(U-B)	Color Temperature (˚K)
upsilon Ori	B0 V	31,500				
lambda Lep	B0.5 IV	26,200				
HD 31726	B1 V	22,600				
eta Aur	B3 V	16,300				
tau Ori	B5 III	13,200				
beta Ori	B8 I	12,200				
mu Ori	A2 V	9,100				
iota Tau	A7 V	8,300				
gamma Lep	F6 V	6,300				
lambda Aur	G2 IV	5,860				
tau Aur	G8 III	4,900				
rho Ori	K0 III	4,500				
alpha Ori	M2 I	3,600				

B-V vs. Temperature
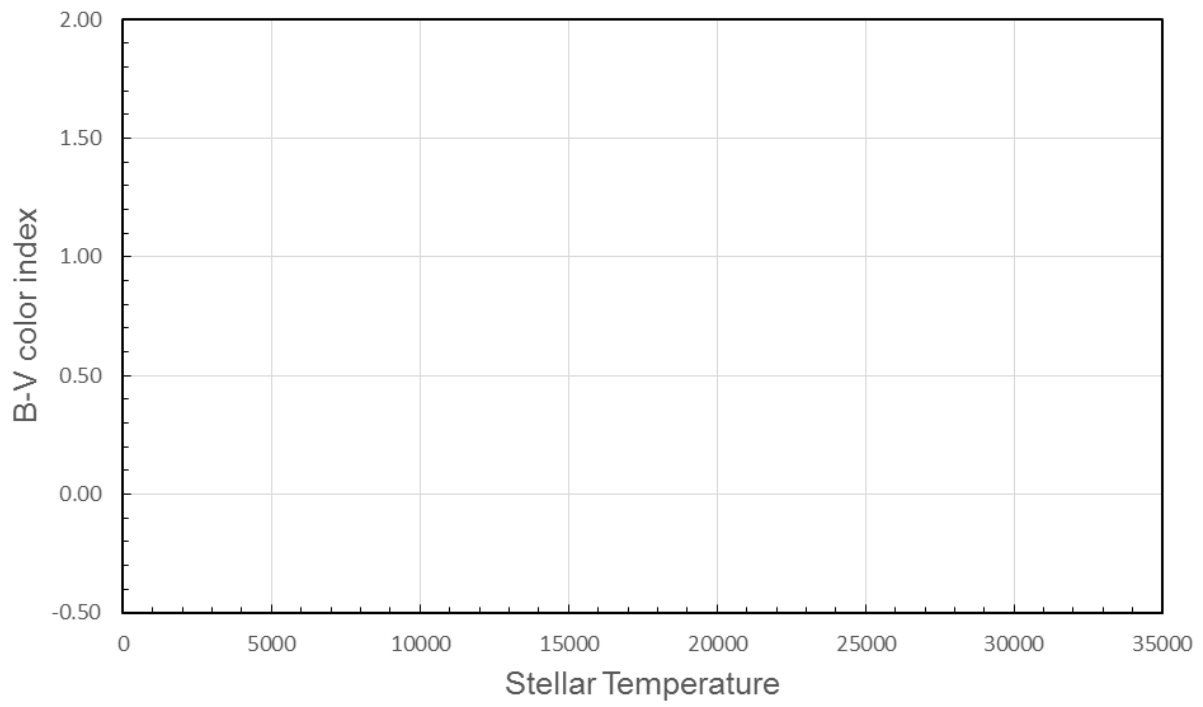
U-B vs. Temperature
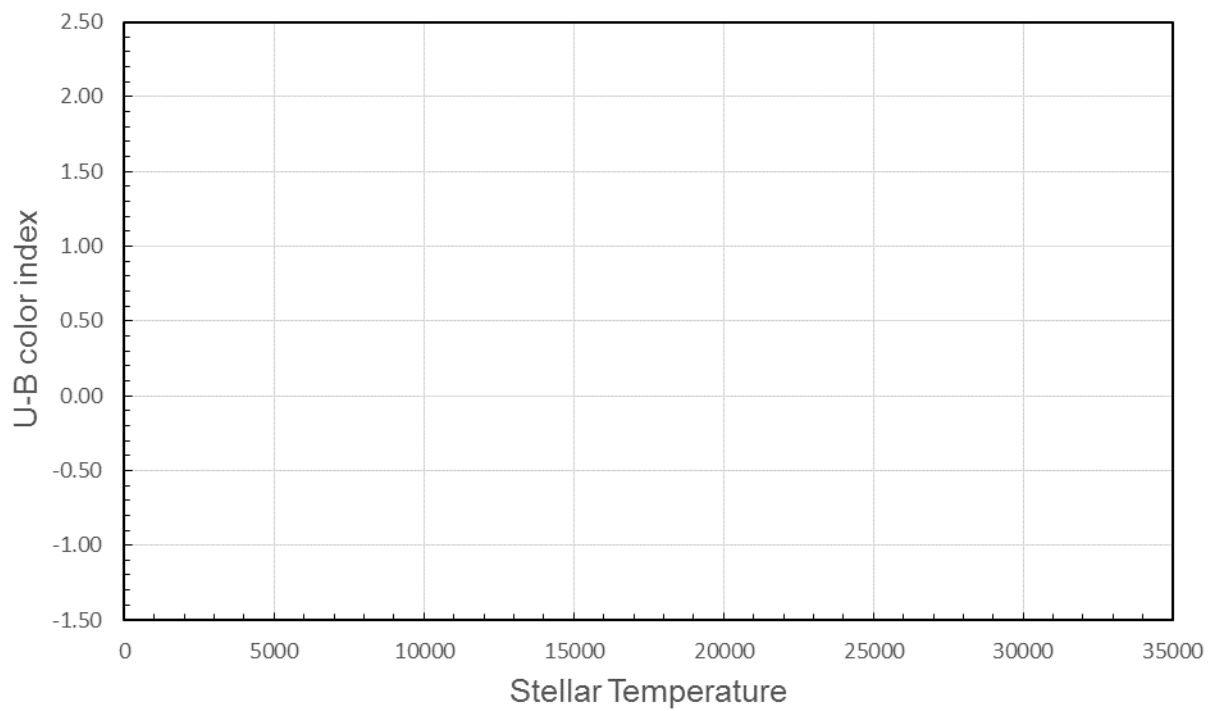
B-V Color Temperature vs. Temperature
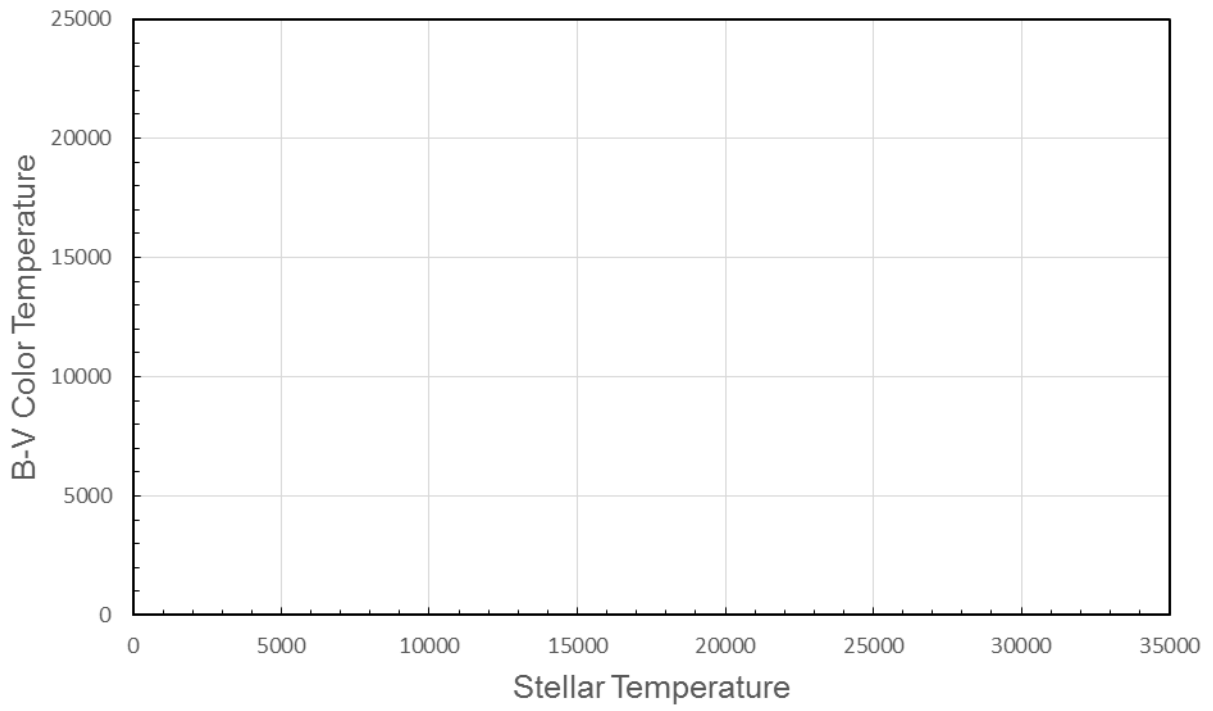
Lab Report Cover Page

Astronomy Laboratory – Stars
Mendel Science Experience 2151
Spring 2022

Lab H
<u>Taking a Star's Temperature</u>

Name: _____

Lab Section: _____

Date: _____

Lab I
Trigonometric Parallax

PURPOSE:

To use simulated observations of the heliocentric parallax of stars to understand the most fundamental method of measuring distances outside the solar system.

EQUIPMENT:

Surveyor's theodolites, computer, scientific calculator, and Villanova's ***Super 'Nova*** computer program
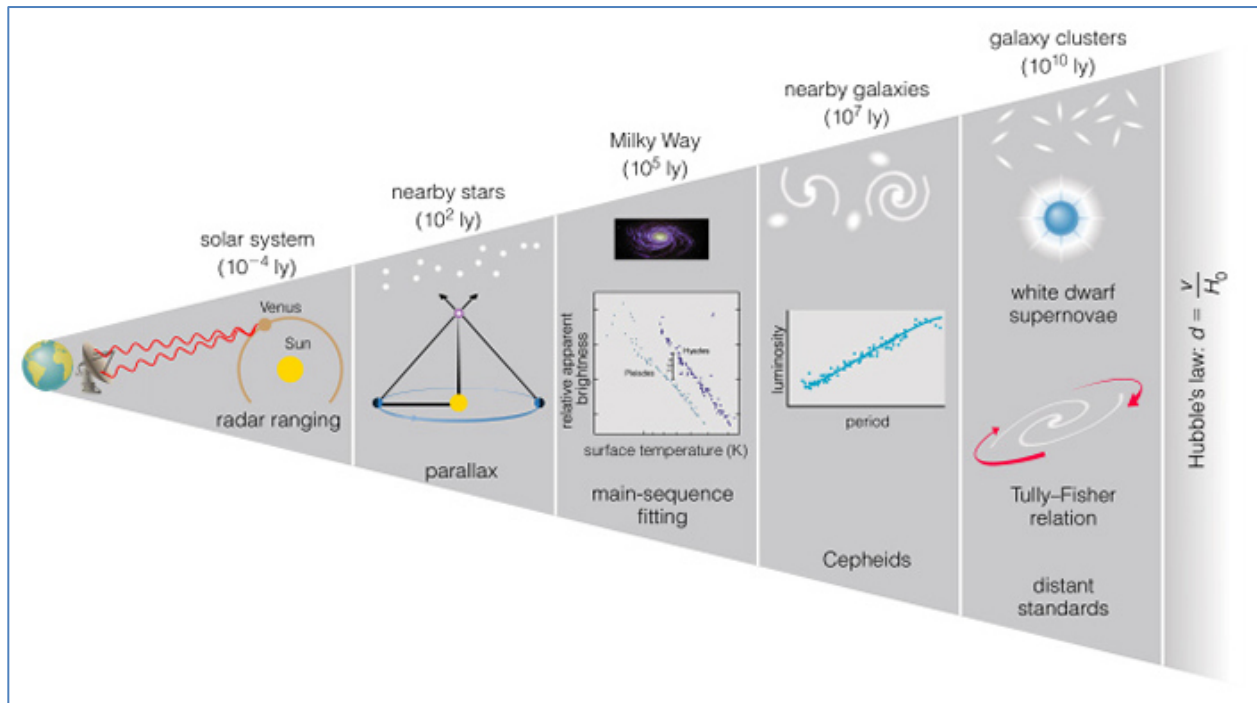
Figure 1: *The Cosmic Distance Ladder*
(Picture credit: Bennett et al. "The Essential Cosmic Perspective")

Introduction

*(NOTE: your instructor may ask you to answer Introductory Questions
1 thru 5 as a pre-lab exercise)*

One of the most important measurements that astronomers make is of the distances to astronomical objects. For example, distances are required for determining many of the fundamental properties of stars and galaxies, such as their luminosities, as well as for determining the size scale of the Universe. For objects on Earth and in the Solar System, it is now possible to use the technique of "radar ranging" to accurately determine distances. In this technique, a beam of radio waves is bounced off a distant object. The time delay between when the beam is sent out and when the reflected beam returns, combined with our knowledge of the speed of light, allows the distance of the object to be computed. The precision of this technique is limited only by our ability to measure the time delays. Beyond the Solar System, however, radar ranging is impossible and astronomers have devised a number of methods that allow distances to ever more distant objects to be estimated. Some of these methods are illustrated in the **Figure 1** on page 1 of the lab. The set of techniques that now allow us to determine distances out to the faintest observable galaxies is known collectively as "The Cosmic Distance Ladder." In general, the techniques for studying the most distant objects depend on calibrations determined by studying closer objects. In fact, the accuracy of the entire Cosmic Distance Ladder ultimately depends on how well we can measure the distance to the relatively nearby stars.

In this lab you will investigate the technique of "Trigonometric Parallax" which astronomers use to measure the distances to the nearest stars. You will see that this is a "fundamental" technique in that it relies on no theoretical assumptions nor on any *a priori* understanding of the properties of stars. You will also see how this technique is similar to a method used by Earth-bound surveyors to measure distance to faraway objects.

THE SURVEYORS METHOD:

Even on Earth, distance measurements are not always simple. Particularly in the days before radar ranging was possible, techniques had to be devised to measure distances to remote or inaccessible objects and locations. One such technique, the "Surveyor's Method," is illustrated in **Figure 2**. The goal is to measure the distance to the tree on the far side of a raging river without getting wet or, possibly, drowning. The technique is straightforward. The first step is to establish a baseline on the accessible side of the river, of known length b. The second step is to measure the two angles β_1 and β_2 using a precise surveyor's instrument called a **theodolite**. Finally, after applying a little trigonometry (i.e., the Law of Sines), the distance to the tree to each theodolite can be calculated from the equations:

$$d_1 = \frac{\sin \beta_2}{\sin \alpha} \times b \tag{1a}$$

and

$$d_2 = \frac{\sin \beta_1}{\sin \alpha} \times b \tag{1b}$$

where α is the apex angle of the triangle in **Figure 2**, and corresponds to the angular size of the baseline as viewed from the distant tree. Since the sum of the interior angles of a triangle is 180°, we have:
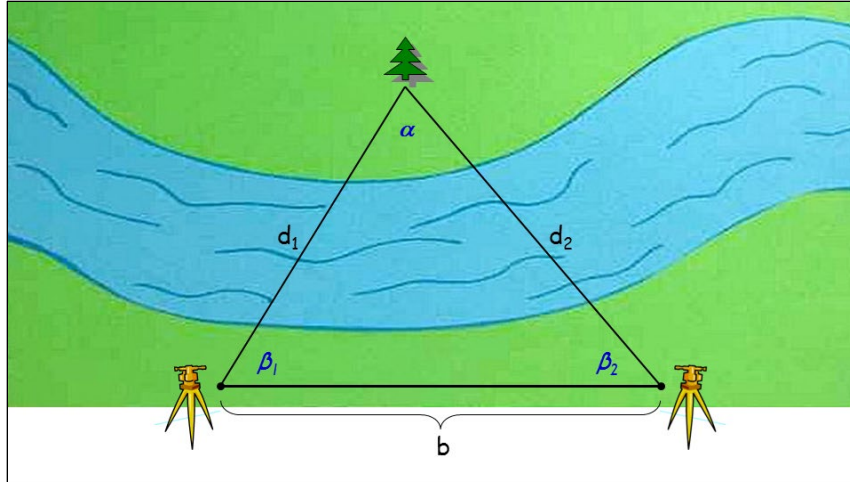
$$\alpha = 180° - \beta_1 - \beta_2 \tag{2}$$

Figure 2: *The Surveyors Method.*

The physical units of the distance in **Equations 1a** and **1b** will be the same as those in which the baseline is measured. The precision of this distance measurement will depend only on the precision of the measurements of b, β_1 and β_2.

Before we discuss the astronomical applications, let's first consider what happens to **Equations 1** when the distance is very large (as compared to the baseline length). From studying **Figure 2**, you can see that, as the distance increases, the angles β_1 and β_2 approach limiting values of 90° and the apex angle α approaches a limiting value of zero. From our knowledge of trigonometry, we know that this also imposes limiting values on the *sine* of these angles. These limits can be summarized as follows:

as d becomes large (i.e., d→∞)	$\beta_1 \rightarrow 90°$ $\beta_1 \rightarrow 90°$ $\alpha \rightarrow 0$	*and*	$\sin\beta_1 \rightarrow 1$ $\sin\beta_1 \rightarrow 1$ $\sin\alpha \rightarrow \alpha$ (radians)

If we substitute these limiting values for the *sine* of the angles into either of **Equations 1a** or **1b**, we get a simplified version of the distance equation which applies for large distances*:*

$$d_1 = d_2 = \frac{1}{\alpha} \times b \qquad\qquad \text{for large } d \qquad (3)$$

where ***the angle α must be expressed in radians*** and the distance will have the same units as the baseline. Thus, for large distances, the distance is inversely proportional to the size of the apex angle α and directly proportional the size of the baseline b. The inverse relationship between the angle α and the distance d can be seen clearly from considering Figure 3.

Later in this lab, you will become surveyors and measure the distance to an object outside the classroom using two theodolites. Because the objects you measure are still relatively close, you will use the more exact version of the distance equation, as given in **Equations 1a** and **1b**.
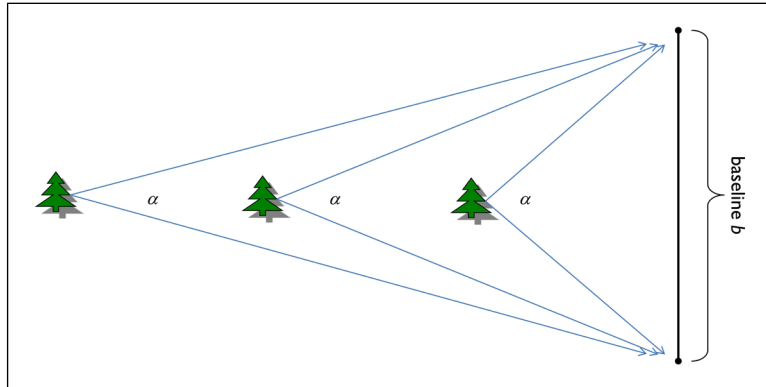
Figure 3: *The inverse relationship between the angle α and distance.*

THE ASTRONOMER'S METHOD:

Astronomers have devised a technique for measuring the distances to nearby stars, which is similar to that employed by Earthly surveyors. It is known as the Trigonometric Parallax method. **Figure 4** shows the general arrangement, which makes these measurements possible. The Earth's orbit around the Sun provides a measurement baseline whose length is well-known, i.e., 2 Astronomical Units (AU). This baseline subtends an angle α as observed from the nearby star whose distance we wish to determine. Due to the large distance of stars, measurements of α and b allow us to determine the distance to the star using Equation (3) and with a precision limited only by the precision of our measurements.

At this point you should be thinking, "*How in the world can the angle α be measured when we can't travel to the star and directly measure it?*"

This is where the "parallax" part of Trigonometric Parallax comes in. "Parallax" is the apparent shift in position of a nearby object with respect to more distant objects when viewed from two different locations. As an example, hold up your thumb at arm's length and look at it first with one eye, and then the other. Your thumb appears to shift back and forth against the background, and this apparent shift is called parallax. It is the result of looking at an object from two different vantage points, that is, from each eye that has a different position relative to your thumb.
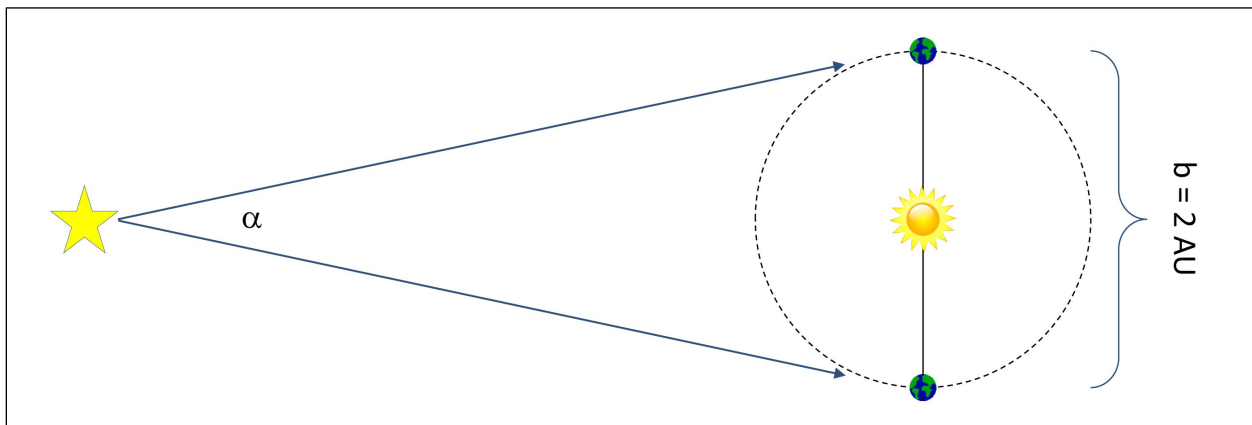

Figure 4: *The Astronomers Method.*

If you move your thumb closer to your face and repeat the experiment you see that the apparent shift of your thumb is greater. Therefore, the more distant an object is the smaller is its shift, or parallax, when viewed from two different positions. For nearby objects, closer than 100 feet or so, your brain measures the parallax by comparing the direction of an object as seen by your two eyes and automatically interprets that parallax as a distance. For this reason, most people find it nearly impossible to touch the tips of their index fingers to each other with one eye closed, but find it easy with both eyes open.

As the Earth moves on its orbit around the Sun, our vantage point for viewing the stars constantly changes. Thus, the relative positions of the stars on the sky actually change during the course of a year, just the way the apparent position of your thumb against a background pattern would change if you moved your head back and forth. Because of the great distances of the stars, however, this change is only measurable for the nearest stars and is far smaller than could be sensed with the naked eye. High quality photographs of the stars, taken at different times of the year, are required to reveal this motion.

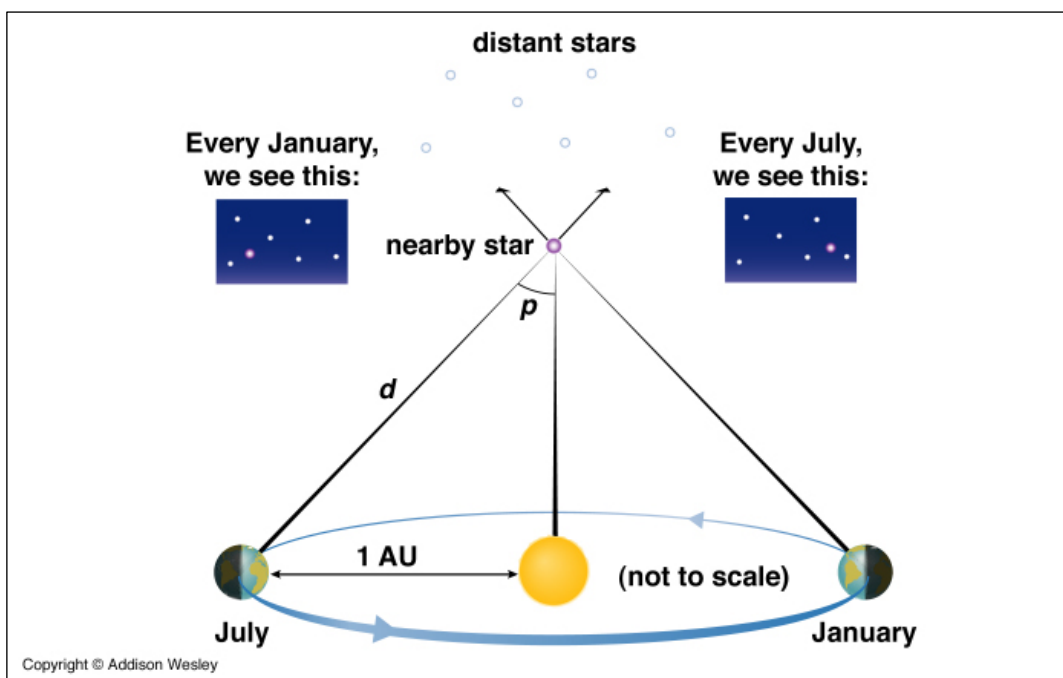The picture below shows how a star's parallax is observed:

Figure 5: Trigonometric Parallax.

Look closely at the two simulated photographs of the stars. You will see that the position of the nearby star has changed relative to the pattern created by the very distant background stars. If you could observe the nearby star continuously over the course of a year, say from one January to the next, you would observe the star to move slowly from one extreme position to the other, and then back again to the original location. Here is the critical part: *the full distance that the star moves on the sky between the January and July observations – which we measure in angular units by comparing the January and July photographs – is the same as the apex angle α in Equation (3) and as shown in Figure 4.* **We don't have to travel to the star to measure α!** Since we already know that the baseline is equal to 2 AU, this measurement allows us to use Equation 3 to determine the distance to the nearby star. The beauty of this technique is that we don't have to know anything about stars to measure their distances. There are no theoretical assumptions in the technique and the precision is limited only by our ability to observe the yearly change in a star's position on the sky, i.e., its parallax.

In practice, astronomers use an even more simplified version of Equation (3) to determine a star's distance. By convention, astronomical parallax angles, represented by the symbol p, are always taken as half the full apex angle (as indicated in **Figure 5**), and are measured in units of arc-seconds rather than radians. Remember that:

1 full circle = 360 degrees (°) = 2π radians
1 degree = 60 arc-minutes (')
1 arc-minute = 60 arc-seconds (")

With these two conventions, and remembering that the baseline distance is 2 AU, Equation (3) reduces to:

$$d\ (AU) = \frac{206265}{p\ (")} \tag{4}$$

where the parallax angle p is measured in units of arc-seconds and the distance d is returned in units of AU. The factor of 206,265 arises because that is the number of arc-seconds in 1 radian.

To further simplify things, astronomers have *defined* a distance equal to 206,265 AU (equivalent to 3.26 light-years) to be one *parsec* (*pc*). With this definition, Equation (4) becomes:

$$d(pc) = \frac{1}{p\ (")} \tag{5}$$

Equation 5 is often known as the Parallax Equation. As you might have guessed, the word *parsec* is a concatenation of the words "parallax" and "arc-seconds". A star with a parallax angle of 1" will have a distance of 1 pc. Another way to look at it is that, from a distance of 1 pc, a baseline of 1 AU will subtend an angle of 1".

INTRODUCTORY QUESTIONS:

Question 1: Suppose the surveyors in **Figure 2** make the following measurements: $b = 50$ feet, $\beta_1 = 87°$, $\beta_2 = 89°$. What are d_1 and d_2? What is the average distance? Show all calculations.

Question 2: What is the critical condition that would allow us to use the simplified form of the distance equation in **Equation 3** rather than the full form in **Equation 1**?

Question 3: What is the distance, in parsecs, of a star with a parallax shift of $p = 0.75$ arc-second? Show all calculations.

Question 4: What is the distance, in light years, of a star with a parallax of $p = 0.75$ arc-second? Show all calculations.

Question 5: What is the distance, in A.U., of a star with a parallax of $p = 0.75$ arc-second? Show all calculations.

Lab Procedure

There are two parts to this lab. In **Part A** you will use the Surveyor's Method and a pair of theodolites to determine the distance to an object outside the classroom. Your instructor will tell you which object to measure. In **Part B**, you will use simulated observations of the night sky to measure the parallax angle *p* and distance *d* to several nearby stars. These two parts can be done in any order.

A. THE DISTANCES TO TERRESTRIAL OBJECTS: THE SURVEYOR's METHOD

There are two surveyor's theodolites set up in the lab next to the window – one at the front of the lab and one at the back of the lab. You will use these to perform a measurement identical to that illustrated in **Figure 2** (except that there is no river). The measurements you will need to make are the baseline size *b* (i.e. the distance between the theodolites) and the two angles β_1 and β_2 between the baseline and the distant object. You will work in pairs or groups, with one student (or group) at the front theodolite and another at the rear theodolite. Your instructor will tell you which distant object to observe.

Step 1: Establish the baseline.

Before beginning, ask your instructor if the theodolites have been leveled for the experiment. If they are not properly adjusted, the measured angles will be incorrect. Once you are assured that the instruments are properly prepared for the meaurement, loosen the vertical and horizontal motion locks to allow the theodolite telescopes to move freely. The locks are indicated by the yellow electrical tape. Turn them counterclockwise to loosen the telescopes and clockwise to lock the telescopes. Point the telescopes so that they are looking directly at each other. Tighten the locks. After the telescopes are locked, small adjustments of the pointing can be made using the fine-adjustment knobs indicated by the blue electrical tape.

After you are satisfied that the two telescopes are pointed directly at each other, press the **green** power button on the theodolite's control panel to activate the display. Next press the "**0SET**" button on the control panel. This will establish the baseline as the zero-angle position of the telescope.

Step 2: Measure the horizontal deflection angle of the distant object.

Once you have "zeroed" the system, loosen the motion locks (yellow tape) and move the telescope so that it is centered on the object whose distance you wish to measure. Center the object in the telescope's crosshairs and lock the telescope. You will have to adjust the telescope's focus to clearly see the distant object. Remember that you can make fine adjustments in the telescope's pointing after it has been locked. **Make sure that both theodolites are pointed to the same part of the object you are measuring!** After you are satisfied that telescope is pointed properly, record the horizontal deflection angle (β_1 for the front theodolite and β_2 for the rear theodolite) in **Table 1** below or on an Excel spreadsheet provided by your instructor. The angle is shown on the instrument's display panel in the **HR** row, in units of degrees, arcminutes, and arc-seconds. If the display reads a negative angle, press the **R/L** button to toggle the telescope's rotational orientation between clockwise and counterclockwise. Your final result should be in the general neighborhood of 80-90°. Share your result with your observing partner. You should both now have measurements for β_1 and β_2.

Step 3: Measure the baseline distance.

Determine the separation between the two theodolites. Use the tape measure provided by your instructor to determine the distance between the two plumb bobs hanging below the theodolites. Make this measurement

as precisely as possible, to a fraction of an inch. Record the baseline distance *b* in Table 1.

Your measurements are now completed. Turn off the theodolites' control panels by pressing the **green** power button again.

Step 4: Compute the distance to the object.

Use your measurements of β_1, β_2 and *b* – along with **Equations 1a** and **1b** – to determine the distances to your chosen object. If you are using Excel for your calculations, remember that it requires angles to be expressed in radians to compute trigonometric functions. When you convert your angles to decimal degrees, you should keep 4 decimal places. Enter your distance result (in feet) in **Table 1**.

> **Question 6:** How well did you do? Your instructor will provide you with the known distance to the object you've observed. Compute the percentage error for your results. Remember that %-error is given by:
>
> $$\% \, Error = \frac{measured \, value - true \, value}{true \, value} \times 100\% \qquad (6)$$
>
> Comment on your results. What distant object did you observe? Were your distances close to the true value? What were the main sources of possible uncertainty in your measurements?

B. THE DISTANCES TO STARS: TRIGONOMETRIC PARALLAX

Step 1: Gathering data

Launch the ***Super 'Nova*** program in the "MSE Astrolab" folder. Choose **Labs**, then **Stellar Parallax** to start the exercise. The computer can simulate five different fields and you will select two to use in this lab. For each field, the computer simulates a photograph of a very small portion of the sky. You can use the **Blink** feature to switch back and forth between three photographs of the same portion of sky taken six months apart and arbitrarily identified as January, July, and January one year later. As you **Blink** between the photographs, the distant stars will not change but the nearby stars will appear to move back and forth just as your thumb appeared to move back and forth as you blinked your eyes at the very beginning of the lab.

Blinking only two plates taken six months apart may show objects that are *really* moving in addition to stars that *appear* to move due to parallax. Objects that are truly in motion are planets, comets, etc. Only additional observations can tell you whether the object is really moving. Pluto was discovered in precisely this way by blinking pairs of plates. Today, new asteroids are discovered with this technique, but they move so rapidly that their motion is apparent on plates taken only a few days apart. Beware! In this lab, there are a couple of objects that are really in motion. **If an object does not move back to its original position one year later, then it is not a star and the apparent motion is due to actual movement.**

To record the data, first write down the number of the field that you have chosen in **Table 2** or on the Excel spreadsheet provided by your instructor. Then blink among the three photographs to identify the stars with measurable parallaxes. Now turn on the overlay Bullseye (click on **Bullseye**), place the Bullseye over one of the stars that exhibits the parallax shift by clicking on it. Make sure the Bullseye is well centered and measure the position of this star for the earlier January date. Record its initial position (*x* and *y*; displayed lower right) in the Table. Click on **Blink** and determine how far this stars' position appears to shift in July. Each ring of the Bullseye counts as one "Grid Unit". **Estimate the shift, *S*, to the nearest 0.1 grid units**

and enter it in the Table. Do this carefully! Do the same for any other star that shows a parallax shift. Note: if you find a moving object (planet, etc.) you will need to measure its coordinates for the first January but leave the shift blank since the movement is actual and not caused by parallax.

After completing your first field, choose another field (**Starfield**, etc.) and repeat the above procedure for each.

Step2: Processing the data

Now that you have identified the stars in your fields with measurable parallaxes and have measured both their positions and parallax shifts, the next step is to calculate their distances. This is done in two parts. First, the shift in grid units must be converted to the parallax angle. Next, this parallax must be converted to a true distance to the star.

The shift, *S*, is actually twice the parallax in grid units. This must be reduced to a real angle, which can be done by knowing the plate scale on the screen. The scale of the screen is 10 grid units = 1 arc-second. Since the parallax is half of the measured angle, you must also divide by a factor of 2:

$$p\ ('') = \frac{S}{10} \times \frac{1}{2} = \frac{S}{20} \tag{7}$$

Using this information, convert the total shift in grid units to the parallax angle in arc-seconds. Record this in your Data Table. Using the information provided in the Introduction (and **Equation 5**), determine the distances (in parsecs) to the stars with measurable parallaxes. Finally, convert these distances to units of light years, using the conversion 1 parsec = 3.26 light years.

> **Question 7:** As you have seen, parallax angles are small. But how small? Imagine someone holding up a dime (diameter = 1.8 cm) at a distance from you such that the dime appeared to have an angular size of 1''. How far away from you would that person be? Show all calculations. (Hint: Use **Equation 3**. In this case, the angular size of the dime is *α* and the dime's physical diameter is the baseline *b*. The angle must be expressed in radians!)
>
> **Question 8:** The diameter of the Sun is 1,392,000 km (i.e., 1.392×10^{11} cm). If you were to construct a scale model of our Galaxy, such that the Sun was shrunk to the size of a marble with a diameter of 1 cm, how far away from the Sun would you have to place a model of the Earth? Express your answer in units of meters. How far away from the Sun would you have to place a model of the nearest star you measured in Part B above? Express your answer in units of km. Show all calculations. (Note, the Earth-Sun distance is 1 Astronomical Unit).
>
> **Question 9:** The fastest spacecraft ever launched by humans is the New Horizons mission, which has recently flown by the dwarf planet Pluto and the Kuiper Belt Object Ultima-Thule. It achieved a velocity of approximately 16 km/sec shortly after launch. If a human-launched spacecraft could sustain such a speed, how long (in years) would it take to reach the nearest star you measured in Part B above? How might such a journey be accomplished? Would you like to make this trip? Show all calculations.

When finished with the computer program, Exit the exercise (**File**, **Exit Parallax Lab**), and then exit the program (**File**, **Exit**). Finally, as usual, shut down the computer completely when finished.

<mark>**NOTE! See the "Useful Units and Conversions" on the next page!!**</mark>

Useful Units and Conversions

1 Astronomical Unit (AU)	= 1.496×10^8 km (Earth-Sun distance)
1 light year	= 63,271 AU
1 light year	= 9.461×10^{12} km
1 parsec	= 3.26 light years
1 year	= 31,557,600 seconds

Table 1: The Surveyor's Method

The object you observed:			
β1 (° ' ")			
β2 (° ' ")			
b (feet)			
β1 (decimal degrees)			
β2 (decimal degrees)			
α (decimal degrees)			
distance d1 (feet)			
distance d2 (feet)			
Average Distance (feet)			

Table 2: Stellar Parallax Measurements

Field #	Star	x position (Jan 1)	y position (Jan 1)	S (grid units)	Parallax (arc-sec)	Distance (parsecs)	Distance (light yrs)
	A						
	B						
	C						
	A						
	B						
	C						

Astronomy Laboratory - Stars
Mendel Science Experience 2151
Spring 2022

Lab I
<u>Trigonometric Parallax</u>

Name: _____

Lab Section: _____

Date: _____

Lab J
The Distance to a Star Cluster

(Adapted from *Contemporary Laboratory Experiences in Astronomy*

PURPOSE:

To learn how astronomers can determine the distance to star clusters using a technique called *Main Sequence Fitting*. You will perform simulated photometry to measure the apparent *B* and *V* magnitudes of 25 stars in the *Pleiades* cluster and use these data to determine the distance to the cluster.

EQUIPMENT:

Computer, *CLEA* computer program ***Photoelectric Photometry***, scientific calculator or Excel spreadsheet.

The Pleiades star cluster.
Photo credit: NASA

Introduction: Absolute Magnitudes and Distance

In an earlier lab, you learned about the technique of *Trigonometric Parallax*, which allows astronomers to measure the distance to relatively nearby stars in a completely assumption-free way. This is important because accurate distances are needed to determine many intrinsic stellar properties. For example, if we know the distance d and measure the apparent brightness b of a star, then we can compute its intrinsic luminosity L from the inverse-square law formula relating these three quantities:

$$b = \frac{L}{4\pi d^2} \tag{1}$$

The luminosity, which is the total energy emitted per second by the star, is related to the rate at which nuclear energy is released deep in the interior of the star. Thus, the distance, while not an intrinsic stellar property itself, plays an important role in uncovering the intrinsic properties. **Equation 1** could also be used to find the distance to a star, if its luminosity were known *a priori*, since the brightness b is always measurable. This method of distance determination is the subject of this lab.

An alternative way that astronomers have of quantifying the luminosity of stars is the *absolute visual magnitude M_V*. It is defined as the apparent visual magnitude V a star <u>would have</u> if it were located at a distance of 10 pc. As with V, which you measured in earlier labs, the more positive the value of M_V, the fainter the star. Because of the way M_V is defined, the difference between a star's actual observed value of V and its intrinsic value of M_V is related to the star's distance. This difference is often called the *distance modulus DM* of a star and is defined as:

$$DM \equiv (V - M_V) \tag{2}$$

The star's distance is related to its distance modulus by:

$$d = 10^{\wedge [(DM+5)/5]} \, pc \tag{3}$$

where the resultant distance is in units of parsecs. Although they look very different, **Equations 1** and **3** are both expressions of the inverse square law of light.

> ***Question 1:*** *Use **Equation 3** to calculate the distance to a star with a distance modulus of zero. Show all steps of your work. Is your answer consistent with the definition of M_V? Explain.*

An important early result from the study of stars with known distances is that, for Main Sequence stars, M_V and spectral type are well correlated. This relationship is shown in **Figure 1** below. The least luminous stars in the diagram are the cool M8 stars on the lower right, and the most luminous are the hot O5 stars on the upper left, with a systematic change in luminosity in between. You should recognize **Figure 1** as a form of the *HR Diagram*, i.e., a plot of a star's luminosity vs. its temperature – but using spectral type and M_V as proxies for temperature and luminosity, respectively.

Because a star's spectral type and *B-V* color are related – we can replace spectral type with *B-V* and re-plot **Figure 1** as a *Color-Magnitude Diagram*. This is shown in **Figure 2**. The beauty of **Figure 2** is that*, if we know that a star belongs to the Main Sequence*, then we need only measure its *B* and *V* magnitudes to determine its distance. The procedure would be to (1) use the measured *B-V* color and **Figure 2** to determine the star's intrinsic M_V, and (2) compute the distance using **Equations 2** and **3** and the observed value of *V*. In this case, only photometric measurements are required, which can be made for large numbers of stars in a relatively short time.
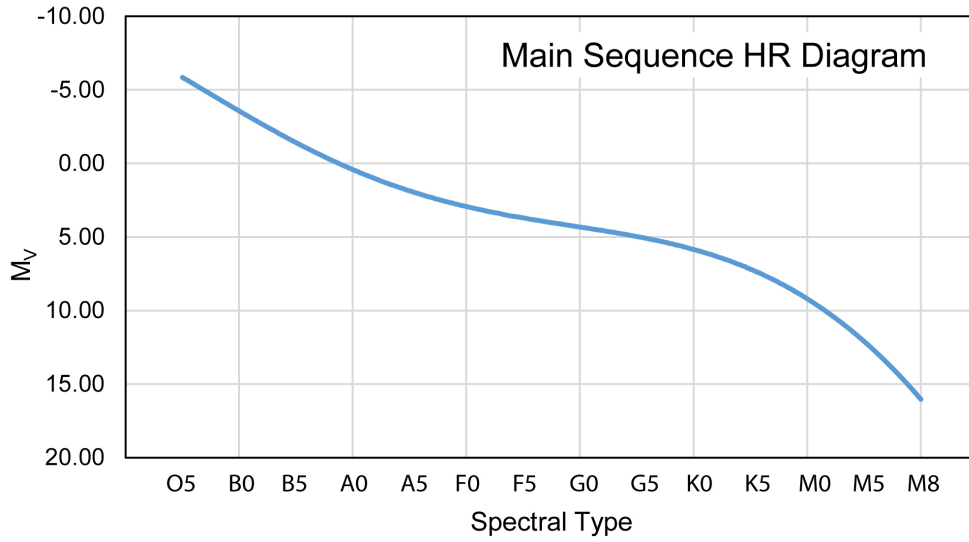
Figure 1: *Absolute Magnitude vs. Spectral Type Relation*
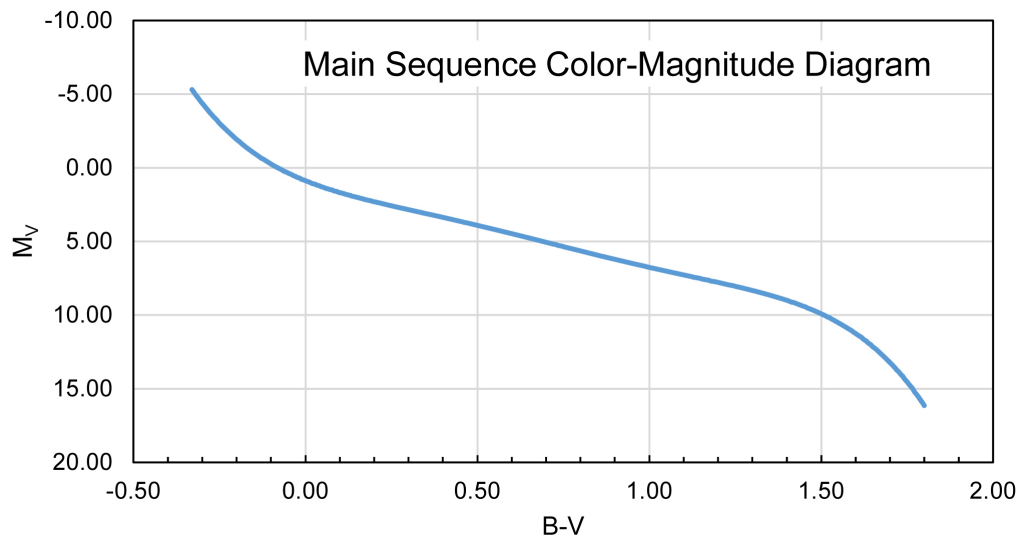

Figure 2: *Absolute Magnitude vs. B-V Color Relation*

The Importance of Star Clusters

The shortcoming with this photometric approach is that the **B-V** color of a star is not sufficient to tell us whether or not the star belongs to the Main Sequence, and some other information would be needed to determine this – such as a carefully determined luminosity class. That's where *star clusters* come in. A cluster is a physically related group of stars which all formed together, and all lie at the same distance from us, thus the distance modulus **DM** is the same for all the stars. If the **V** magnitudes and **B-V** colors of cluster stars are plotted against each other, the Main Sequence stars will form a pattern like that seen in **Figure 2** and thus be readily identifiable. Other types of stars, such as the intrinsically more luminous giants and supergiants will lie above the main sequence line and also be clearly distinguished. Once the cluster's Main Sequence stars have been identified, the calibrated relation in **Figure 2** can be used to determine the stars' M_V values and distance moduli. The distance to the cluster can then be computed as the average of the individual stellar distances, which should all yield consistent results. This general technique is known as the *Main Sequence Fitting* technique.

Main Sequence Fitting can be applied to star clusters at far greater distances than feasible for *Trigonometric Parallax*. Knowing the distances to star clusters in our Milky Way galaxy, and thus their 3-dimentional spatial distribution, is important for understanding the structure of the Milky Way. An added value of determining a cluster's distance from its Main Sequence stars is that, once the distance is known, the luminosities of all its other stars can be computed. In this way we can discover the intrinsic properties of other, less well-understood, types of stars.

Lab Procedure

In this lab you will perform the *Main Sequence Fitting* technique on a relatively nearby group of stars called the *Pleiades star cluster*. This cluster – shown on the first page of the lab – is in the constellation *Taurus* and is readily visible to the naked eye. You will first perform simulated photometry of the cluster's stars, and then determine the mean distance modulus.

You will use a realistic simulation of a *photoelectric photometer* attached to a moderate-sized research telescope to collect data on 25 stars in the region of the *Pleiades* star cluster. The simulation allows you to select different filters to isolate the wavelength regions of interest and to select your desired integration time. For this lab, you will measure the **V** and **B** magnitudes, and determine the **B-V** color for each of your program stars. To save you time, the simulation will automatically compensate for the effects of atmospheric extinction. You, however, will have to measure the brightness of the sky, so that its effects can be removed from your data.

A. MAKING THE MEASUREMENTS

Step 1: Launch the simulation

In the "MSE *Astrolab"* folder, launch the *Photoelectric Photometry* program. Click **File, Log In…** to open the Student Accounting screen. Enter your name, without using any punctuation marks, then click **OK** to continue.
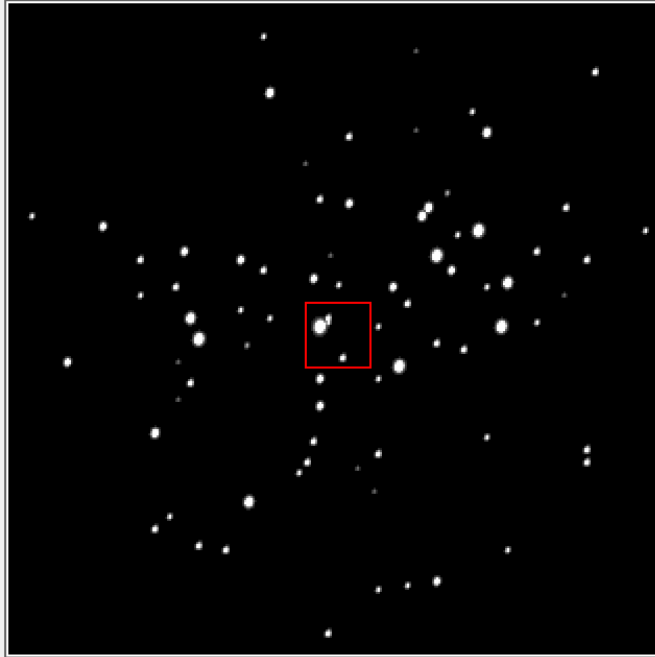
Figure 3: *The Pleiades cluster as seen through the finder telescope in the simulated observatory.*

Step 2: Open the observatory

Click ***File, Run*** to activate the observatory, and then take a moment to study the various controls available to you. You now have command of a 16" research-grade astronomical telescope equipped with a photoelectric photometer. Click on **Dome** to slide the observatory dome doors open, which will reveal the night sky in the direction of the *Pleiades* star cluster, as seen through the finder telescope. This view of the cluster is shown above in **Figure 3**. The region shown is the same as illustrated in the photo on page 1 of the lab. Compare the two to verify this. You can use **Figure 3** to keep track of the stars you observe for your distance determination. Note that the brightness of the stars in **Figure 3** (and in the simulation) is indicated by the symbol size – the brighter the star, the larger the symbol.

You should observe that the stars visible through the finder telescope seem to be drifting towards the right, which is the westward direction. This is because the Earth is rotating and, in general, causes astronomical objects like the Sun to "rise" in the East and "set" in the west. Telescopes are equipped with a motor drive which moves the telescope in the same direction as the drift and at the same rate. The motor (often called the clock drive) cancels the effect of the Earth's rotation and the stars seems to stand still, permitting extended study. Turn on **Tracking** and note how the stars cease to drift.

Step 3: Move to a star

Once the dome is opened, a small red square in the center of the view appears (see **Figure 3**). This is the field of view seen by the photometer, when activated. The star you wish to observe must be carefully positioned within this field to be measured. To move a star into the field, click and hold one of the **N**, **S**, **E**, or **W** buttons. Note that these directional controls move the telescope, not the sky! Moving the <u>telescope</u> to the west appears to make the <u>stars</u> move in the opposite direction. Try it. The **Slew Rate** control changes how much the telescope is moved each time the directional controls are used. Try it, too.

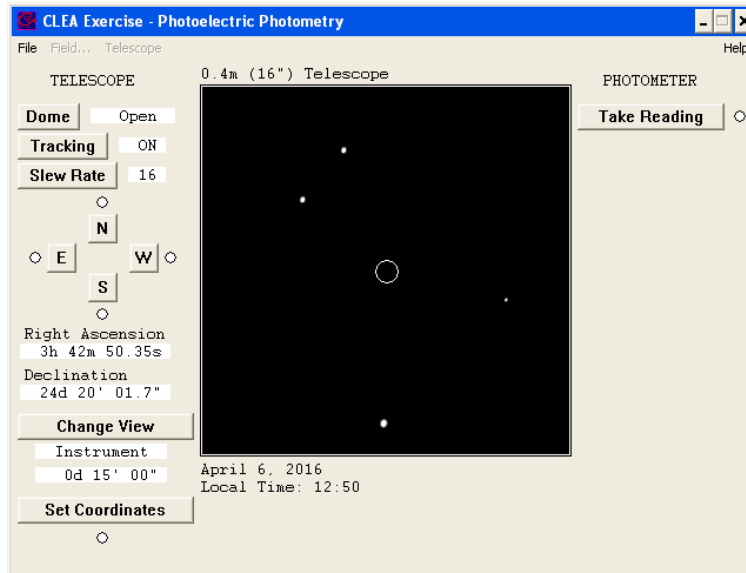Now move the telescope so that one or more stars appears inside the field of view of the photometer.

Figure 4: *The "Instrument Mode" screen*

Step 4: Take sky readings

Click on **Change View** to enter *Instrument* mode and activate the photometer. The computer screen will look as shown in **Figure 4**. The aperture, through which light passes to the photometer, is shown as the small circle in the center of the field. Because the aperture is much larger than the star under study, and the sky is not perfectly dark, the sky within the aperture contributes a certain number of photons to the measurements of the stars. These unwanted photons are counted by taking a "sky" reading, with no star in the aperture. The star to be measured is then centered in the aperture and a "star+sky" reading is taken. Then the true photon count for the star by itself is approximated by the "star+sky" reading minus the "sky" reading. It is important that both the sky and the star reading be taken through the same color filter and with the same aperture size.

Move the telescope until the aperture (circle) is free of any star, then click on **Take Reading**. A new window will open ("CLEA Photometer Window"), containing the photometer controls. Clicking the **Filter** button cycles through the *U*, *B*, and *V* filters. Make sure the integration time is set to 10.0 seconds (by clicking the **Seconds** button) and that the number of integrations to be performed is set to 3 (by clicking the **Integrations** button). Now, whenever the **Start Count** button is clicked, the photometer will count photons for a total of 30 seconds, in three groups of 10 seconds each.

Select the *V* filter and press **Start Count.** When the measurement is finished, record the result (Mean Sky Counts/Sec) in **Table 1** (or in an Excel spreadsheet if provided by your instructor). Now switch to the *B* filter and repeat. When you are finished, press the **Return** button to close the CLEA Photometer Window. The observatory computer recognized that these were blank sky measurements by the absence of a star in the aperture. Every stellar measurement made from now on will have this sky brightness removed before the final magnitude is returned. In a real observing session, an astronomer would likely measure the sky brightness in a small patch of blank sky adjacent to each target star, to account for spatial and temporal variability. To save time, however, we will assume that the sky brightness is spatially uniform and unchanging throughout the night. Thus, this single set of sky measurement will be applied to all the stellar measurements.

Step 5: Measure the stellar B and V magnitudes

Move the telescope to center a star in the aperture. The right ascension and declination of the aperture center are displayed in the lower left of the screen. Click **Take Reading** and select a filter (***B*** or ***V***) and an integration time in the Photometer Window. Use short integration times for bright stars, and long integrations for faint stars.

Now click **Start Count**. The computer will take a series of three integrations and display the individual and average photon counts in the Photometer Window. After the integrations are completed, the *apparent magnitude* of the object under study in the chosen filter is displayed in the **Magnitude** box (lower right corner). This value has been corrected for the effect of sky brightness. Also displayed is the signal-to-noise ratio (**S/N Ratio**) of the reading. A high S/N means you have a lot of desired photons, and only a little noise. You should strive for S/N figures of at least 100. You can increase the S/N by increasing the integration time or increasing the number of integrations. For most of your measurements, an integration time of 1 second will be sufficient. If the resultant S/N value is too low, then switch to 10 sec. (Never use 100 sec - it takes too long!) Always use three integrations.

For each of 25 stars in the *Pleiades* star cluster, record the Star Name, ***B*** magnitude, and ***V*** magnitude in **Table 2** (or in an Excel spreadsheet).   Record all magnitudes to the nearest 0.01 magnitude.   Be sure to include both bright and faint stars in your sample to cover the Main Sequence over the widest possible range.

<mark>NOTE: You must observe at least two stars with V > 12.</mark>

B. PREPARING THE DATA

Step 1: Compute the B-V color index

Compute the ***B-V*** color for each star by subtracting its ***V*** magnitude value from its ***B*** magnitude. Record the result in **Table 2** (or in your spreadsheet). Remember from an earlier lab that hot blue stars have low, and even negative, ***B-V*** values while cooler red stars have ***B-V*** values of +1 or greater.

Step 2: Make a Pleiades color-magnitude diagram

Using the graph provided at the back of the lab (or your Excel spreadsheet), make a color-magnitude diagram of your *Pleiades* data, using the ***V*** magnitude on the y-axis and the ***B-V*** color on the x-axis. If you are using Excel, your plot should be modeled on the blank graph in the back of the lab. **Be sure that the y-axis is oriented so that the brighter stars are at the top of the diagram.**

Step 3: Make a calibrated Main Sequence color-magnitude diagram

Now plot, in the same fashion as above (by hand on the graph provided or with Excel), a calibrated color-magnitude diagram for Main Sequence stars using the data in **Table 3**, which is taken from Allen's *Astrophysical Quantities,* 4th Edition (2000). This version of the diagram, which you have already seen in **Figure 2**, uses the luminosity of the stars (in the form of M_V) rather than the apparent brightness (in the form of ***V***) on the y-axis. Your graph should resemble **Figure 2**.

> ***Question 2:*** *Do you recognize the presence of the Main Sequence in your V vs. B-V diagram for the Pleiades stars? Describe it. Do all your stars seem to be Main sequence stars? Identify the Main Sequence on your plot. (If you are using Excel, you may have to wait until you print*

out the plot at the end of lab and then use pencil to do the labeling.)

C. DETERMINING THE PLEIADES DISTANCE

Step 1: Measure the distance modulus DM

If your instructor provided you with an Excel spreadsheet for recording your data, follow his/her directions for determining the distance modulus **DM**. Record your value for **DM** in **Table 4** and then jump down to **Question 3**.

If you used the graphs provided with the lab and plotted your data by hand, follow these instructions:

> *Overlay and slide the two graphs up and down until the Main Sequence on the Absolute Magnitude plot best aligns with the Main Sequence on your Apparent Magnitude (Pleiades Cluster) graph. Keep the y-axes precisely parallel. Seek a best fit for the central portion of the combined patterns. The cool red stars in the lower right of your cluster graph are quite scattered and may not fit very well.*

> *Consider what you are doing: You have graphs of two groups of main sequence stars. One graph is in terms of visual Apparent Magnitude (V) and the other one in visual Absolute Magnitude (M_V). When the patterns are matched, each star on the combined main sequence can be described either in terms of V or M_V. It just depends on which scale you read.*

> *So, once the two Main Sequence patterns are aligned, a fixed relationship is established between the Apparent and Absolute Magnitude scales, no matter where you read the y-axis or which star you pick. So, pick <u>any</u> convenient magnitude on the Absolute Magnitude scale and read its corresponding Apparent Magnitude on the paper scale. **Read each scale to the nearest 0.1 magnitude**.*

> *You must sketch the Main Sequence curve from the Absolute Magnitude plot onto your Pleiades graph.*

Now that you've determined the distance modulus **DM,** answer the following question:

> **Question 3:** *What distance modulus (DM) do you find? If you used the two graphs provided with the lab, then list the values of the Absolute Magnitude and the corresponding Apparent Magnitude which lead to this determination.*

Step 2: Compute the distance

Use **Equation 3** to compute the distance to the Pleiades in parsecs. Then convert to light-years (where 1 pc = 3.26 ly). Record these values in **Table 4**.

D. SUMMARY QUESTIONS

Question 4: *How well did you do? Recently (in 2014), astronomers at the National Radio Astronomy Observatory determined a precise distance to the Pleiades cluster of 443 light years, with an estimated uncertainty of only about 1%. Assuming that this is the correct result, what is the percentage error in your measurement? What do you think are the major sources of uncertainties in your result?*

Compute the % Error of your measurement using the following formula:

$$\% \text{ Error} = \left(\frac{d_{measured} - d_{accepted}}{d_{accepted}} \right) \times 100 \tag{4}$$

Enter this result in **Table 4**.

Question 5: *Look over your Pleiades color-magnitude diagram and the stars you identified as main sequence stars. Does your sample contain any red giant stars? How would you recognize red giants on your plot? How are they different from main sequence stars of the same temperature? If you do have red giants, list them in **Table 5** and label them on your plot.*

Question 6: *How would your distance estimate for the Pleiades have changed if you had based your distance modulus measurement only on a single red giant star that you mistakenly thought was a main sequence star? Explain your answer.*

Question 7: *Consider a star in the direction of the Pleiades with **V** = 18.0 and **B-V** = 0.1. It would seem curiously out of place with respect to the main sequence. Have you talked in class about any type of star that might lie in this place on the color-magnitude diagram? (Yes, you probably have.) How would such a star be different from main sequence stars of the same temperature?*

*** EXTRA CREDIT QUESTION***
Note: because this is extra credit, you may not ask your instructor for help!

Question 8: *If you lived out in the Pleiades and looked in the direction of Earth, how bright would the Sun be in the nighttime sky? I.e., what would its **V** magnitude be? Would the Sun be visible to the naked eye? (Our eyes can see stars as faint as **V**~ 6.) Explain your reasoning.*
*Note: The Sun is a G2 V star with **B-V** = 0.62.*

Table 1: Sky Brightness

Filter	Mean Sky (counts/sec)
B	
V	

Table 2: Pleiades Photometry Data Sheet

	Star Name	*B*	*V*	*B-V*
1				
2				
3				
4				
5				
6				
7				
8				
9				
10				
11				
12				
13				
14				
15				
16				
17				
18				
19				
20				
21				
22				
23				
24				
25				

Table 3: Main Sequence Color-Magnitude-Spectral Type Calibration

Spectral Type	Absolute Magnitude	B-V
O5	-5.70	-0.33
B0	-4.00	-0.30
B5	-1.20	-0.17
A0	+0.65	-0.02
A5	1.95	+0.15
F0	2.70	0.30
F5	3.50	0.44
G0	4.40	0.58
G5	5.10	0.68
K0	5.90	0.81
K5	7.35	1.15
M0	8.80	1.40
M5	12.3	1.64
M8	16.0	1.80

Table 4: The Distance to the Pleiades

DM	d (pc)	d (light years)	% error

Table 5: Possible Red Giant Stars

	Star Name
1	
2	
3	

Pleiades Color-Magnitude Diagram
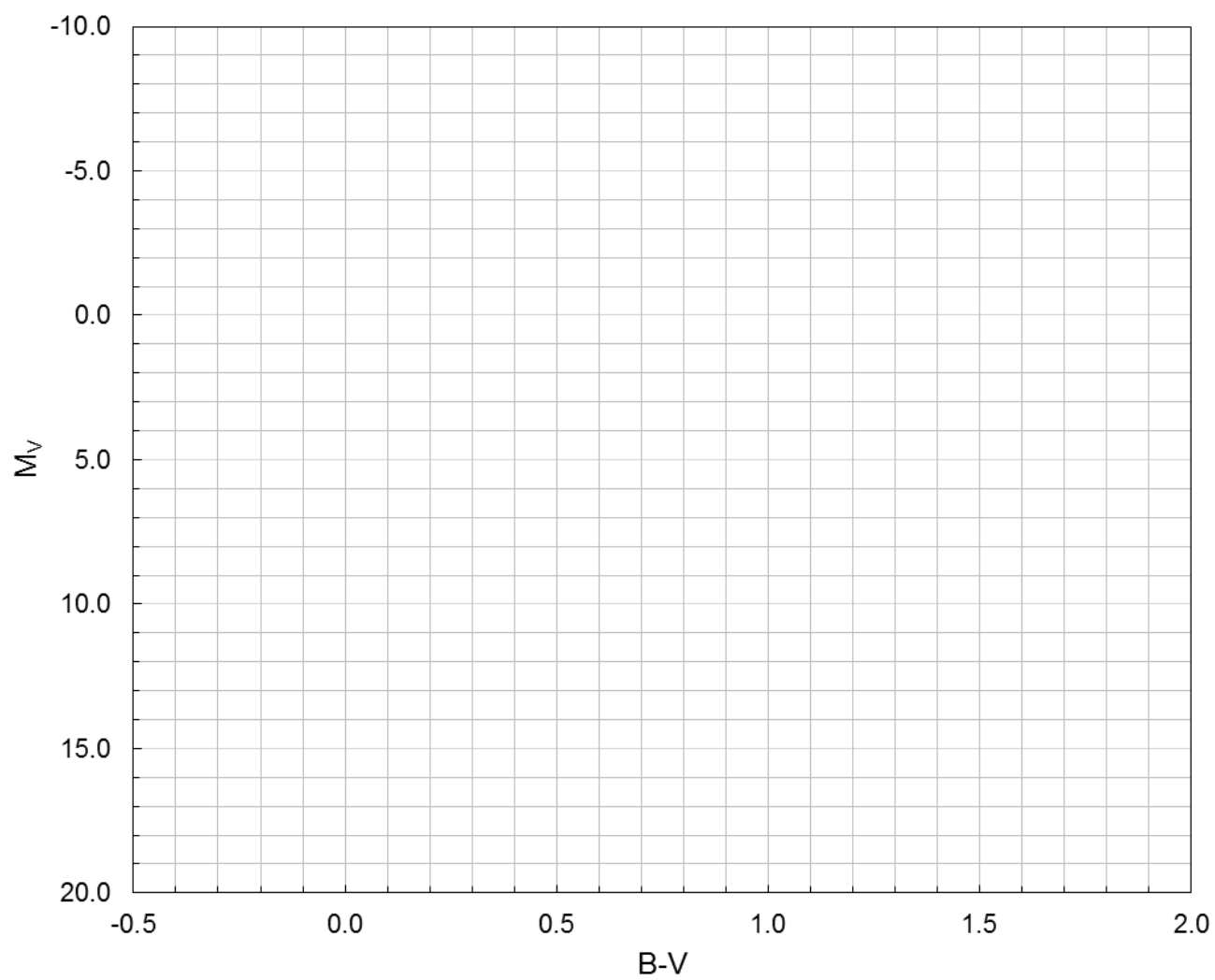
Empty grid plot with y-axis labeled V ranging from -5.0 at top to 25.0 at bottom (increments of 5.0) and x-axis labeled B-V ranging from -0.5 to 2.0 (increments of 0.5).

Main Sequence Color-Magnitude Diagram

Lab Report Cover Page

Spring 2022
Mendel Science Experience 2151
Astronomy Laboratory - Stars

Lab J
<u>Photoelectric Photometry of the Pleiades</u>

Name: _____

Lab Section: _____

Date: _____

Lab K
Measuring the Mass of the Andromeda Galaxy

(Adapted from *Astro Labs*, University of Maryland)

PURPOSE:

To measure the mass of the *Andromeda Galaxy* using the orbital motions of the stars and gas clouds in its disk. The distribution of mass within the visible galaxy will be examined and evidence found for significant amounts of "unseen matter."

EQUIPMENT:

AstroLab computer program, ruler, and Excel spreadsheet or calculator.

The Andromeda Galaxy
Photo credit: Robert Gendler 2002

Introduction: Galactic Rotation & Mass

One of the most fundamental properties of a galaxy is its mass, which plays a critical role in galaxy formation and evolution. The masses of astronomical objects are, in general, difficult to measure because we are not able to put the objects on a scale or a spring balance. Astronomers have had to be clever and figure out what kinds of observations – necessarily made from distances of millions, billions or trillions of miles – can shed light on an object's mass. This is where gravity comes in.   The force of gravity between any pair of objects is known to depend on both the masses of the objects and the separation between them. And so, if we can observe the gravitational interaction between objects in space, we can work backwards and determine the masses of the objects. This is the primary technique used for determining the masses of individual stars, namely, through measurements of the orbits of *binary stars* (which are stars gravitationally bound to each other and orbiting a common center).

Gravity can also be used to measure the masses of entire galaxies.   The most visible parts of *spiral galaxies* are large, flattened, rotating disks consisting of stars and interstellar clouds of gas and dust particles.   These move within the disk in nearly circular orbits around the centers of their galaxies, and their orbital properties are controlled by the collective mass of the galaxy itself.   Thus, by measuring the orbital properties of the components of a galaxy – specifically, the orbital radii and speeds – the mass of the galaxy can be determined.

The key to deriving a spiral galaxy's mass is *Kepler's Third Law,* which was discovered by the German mathematician/astronomer Johannes Kepler in the early 1600's from observations of Solar System objects. *Kepler's Third Law* can be stated as:

$$M_{<a} = K \, \frac{a^3}{P^2} \tag{1}$$

where, in our application, a is the radius of a star's or gas cloud's orbit around its galactic center, P is the orbital period, $M_{<a}$ is the mass of the galaxy within a distance a from the center, and K is a constant whose value depends on the units chosen for M, a, and P.   Measurements of a and P allow the enclosed mass to be determined.   If these measurements are made at a range of orbital radii, then we can build up a picture of how the mass within a galaxy is distributed, as well as an estimate of the galaxy's total mass.

This is the procedure you will carry out in today's lab, in which you will make measurements of a nearby spiral galaxy called *Andromeda*. The *Andromeda* galaxy, shown on the first page of the lab, is usually considered to be a twin to our own Milky Way galaxy and shows us what our galaxy would look like if viewed from a distance of about 2.5 million light years. You will investigate the rotation properties of *Andromeda* by measuring the orbital radii and speeds of its stars and interstellar clouds at a variety of radii and then determine the total mass and mass distribution of *Andromeda*.   This will lead to a startling discovery about the nature of the matter making up galaxies in our Universe.

Lab Procedure

Measuring the orbital speed of stars or interstellar clouds in other galaxies is difficult because we do not have the luxury of simply watching a galaxy rotate and determining the orbital speeds from the time needed to complete a rotation. This would generally take hundreds-of-millions of years, which is beyond the attention span of most people. Indeed, because of the great sizes and distances of galaxies (and their relatively slow rotation speeds), it is never possible to directly observe any *transverse motion* – i.e., motion across the line-of-sight – in a galaxy's components. The best we can do is to measure the *radial motion* – i.e., motion along the line-of-sight – of objects in a galaxy using the *Doppler Effect*, which you learned about in your lecture class. By carefully choosing the locations of the observations, and knowing something about the orientation of a galaxy, these radial motions can be corrected to yield the true orbital speed of a star or interstellar cloud. From this, the mass of a galaxy and its mass distribution can be inferred.

A. TILTED ROTATING DISKS

Step 1: Launch the Simulation

Launch the *__AstroLab__* program in the "MSE Astrolab" folder and choose **Galactic** on the opening screen to start the lab. The first part of the program demonstrates the motion of a rigid circular disc, like a Frisbee or a DVD. The main part of the display is a representation of the disc with arrows indicating the direction of motion. To the right is a "side view" showing the observer's eye looking at the rotating disc. As you move the cursor around on the disc with the mouse, the displays show the distance from the center of the disc, the true rotational velocity at that point, and the line-of-sight velocity that would be measured via the Doppler Effect.

Step 2: Investigate the effects of inclination

In this first display, the disc is being viewed face-on, corresponding to an inclination angle of $i = 90°$. In this case, all of the rotational motion is in the transverse direction (i.e., in the plane of the sky) and none of it is in the radial direction. Verify this by moving the cursor around in the disk. You should see that the "Velocity in line of sight" (i.e., the radial component which would be measured by the Doppler shift) is always 0.

Now cycle through a set of different inclination angles by clicking **ChangeInclination.** Notice that the apparent shape of the disk becomes flatter as the inclination approaches zero, although the largest diameter, called the *major axis*, is always the apparent diameter of the true circular form. If the true shape of the disk were known to be circular, then it would be possible to determine the inclination from the apparent flatness of the projected disk. You will do this below in Step 3 to measure the inclination of the *Andromeda* galaxy.

Now let's see how the inclination angle affects the line-of-sight velocities. <u>Remember, these are the velocities you would actually measure using the Doppler shift.</u> Select one of the inclinations angles between 0° and 90° (i.e., either, 18°, 22.5°, 30°, or 45°). Put the cursor at the position X = –450, Y = 0 and record the values of the "Rotational Velocity" and "Velocity in line of sight" in the appropriate places in **Table 1**. Now do the same for the rest of the X, Y positions listed in **Table 1**. Notice that these positions sample the rotation across the major axis of the disk, on both sides of the center.

Now look at **Figure 1** below. It is a "side view" of the rotating disk (or galaxy), like that seen in the upper right of the simulation. **Figure 1** shows that, when the line-of-sight velocities v_{los} are measured along the major axis, there is a simple trigonometric relationship between v_{los} (which we can measure) and the actual rotational speed v_{rot} (which we want to know). This relationship is:

$$ABS(v_{los}) = v_{rot} \times \cos i \tag{2}$$

Solving for v_{rot} yields:

$$v_{rot} = ABS(v_{los}) \div \cos i \tag{3}$$

You will recall that v_{los} can be either negative or positive, depending on whether the object being observed is approaching or receding, respectively. **Equation 2** recovers only the magnitude of v_{los}, not the direction – hence the use of the absolute value function. Let's test **Equation 3** using the data you just collected. For each of the v_{los} measurements you entered in **Table 1**, compute the quantity $ABS(v_{los}) \div \cos i$ and enter in the appropriate column. If our understanding is correct, then this modified value of v_{los} should be identical to the actual value of v_{rot} you measured. If you are using Excel, remember that angles have to be converted from degrees to radians before performing trigonometric calculations.

> **Question 1:** Do your values for $ABS(v_{los}) \div \cos i$ agree with the true rotation speeds v_{rot} for all positions along the major axis? Do you agree that, if you can measure the line-of-sight velocities of objects along the major axis in a tilted spiral galaxy, then you can determine their true orbital velocities (provided you know the inclination angle i)?

Step 3: Measure the inclination of the Andromeda galaxy

Since the distributions of stars and interstellar clouds within spiral galaxies are known to resemble thin circular disks, the flattened shape of the *Andromeda* galaxy shows that it is not being viewed in a face-on orientation (i.e., $i \neq 90°$). To determine the inclination angle *i,* we simply need to assume that the true face on shape is circular and then measure the relative sizes of the major axis and the *minor axis.* The minor axis is the smallest diameter of the oval-shaped image and is perpendicular to the major axis, passing thru the apparent center of the galaxy. (In the simulation, it is the line connecting points A and B.)
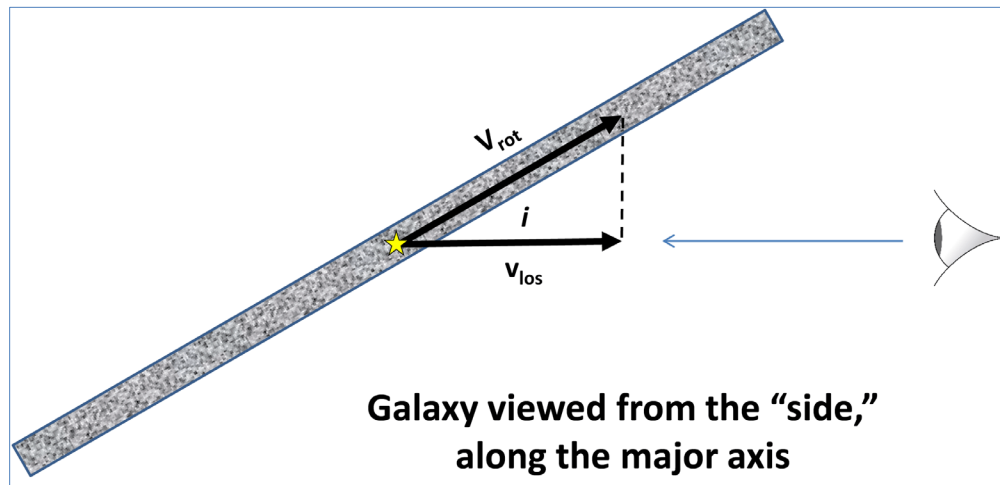

Figure 1: *The effect of inclination on observed velocities*

Using the photograph of *Andromeda* supplied on the last page of the lab and a ruler, measure the long diameter and the short diameter of the *Andromeda* image. You will have to be consistent in your estimate of the locations of these diameters and the extent of the galaxy in both directions. Enter the results in first two lines of **Table 2**.

Now refer back to **Figure 1**. A little trigonometric insight should show you that the relationship between the major and minor axis lengths is given by:

$$\text{minor axis} = \text{major axis} \times \sin i \qquad (4)$$

This relationship can be rearranged to express the inclination in terms of the major and minor axis lengths:

$$i = \sin^{-1}\left(\frac{\text{minor axis}}{\text{major axis}}\right) \qquad (5)$$

Compute Andromeda's inclination angle *i* from your measurements and enter it in both radians and degrees in the third and fourth rows of **Table 2**. If you are using Excel, remember that the inverse sine function is given by "ASIN". Also note that Excel will return the angle *i* in radians. To convert to degrees, use the "DEGREES" function.

B. THE ROTATION CURVE OF ANDROMEDA

Click **NextPart** on the lab menu bar to proceed to the measurement of *Andromeda's* rotation properties.

Step 1: Set the major axis

The computer screen shows a picture of the Andromeda Galaxy. Above the image there is a line with two small filled circles at both ends and another to indicate the center. Your first job is to move this line into a position so that it represents the major axis of *Andromeda*. You can control the tilt of the line and the location of its center by clicking and dragging the filled circles. Now, position the line so that it runs down the long axis of the galaxy with the center circle at the center of the galaxy.

Step 2: Measure the Doppler shifts

If we were actually doing this experiment, we would take spectra at various points along the major axis line, measure the Doppler shift and determine the radial velocity. We would then measure on the photograph the distance from the center to the various points at which the spectra were taken. These measurements would usually be made in some convenient units (e.g., mm) and then converted into real distances (e.g., in parsec) by using the known distance to *Andromeda*. This is obviously a time-consuming task and, to save time, all these calculations will be done for you by the simulation program. Notice that there is a small "bead" that moves along the line as you move the cursor around. This bead marks the point where you are "taking a spectrum" and measuring the radial velocity via the Doppler shift. The readouts at the bottom of the screen give the line of sight (radial) velocity of the point at which the bead is located and its actual distance from the center of the galaxy ("*a*") in units of kiloparsecs (kpc), where 1 kpc = 1000 pc.

Starting at the north-east (upper left) corner of *Andromeda*, position the cursor at each of the *a* values listed in the first column of **Table 3** and record the line-of-sight velocity, v_{los} in column 2 of the table. Work

down thru the center of Andromeda and end up in the south-west corner, 27 kpc from the center.

Step 3:   Determine the systemic motion

If you recall the demonstration of the rotating disks (Part A of the lab), you should realize that there is something wrong. For the rotating disk, the center point did not move and had a zero velocity. However, the radial velocity at the center of this galaxy is far from zero.   This is because there are other motions involved besides simple rotation. *Andromeda* and our galaxy may be getting closer together or further away from one another. Our Sun is presumably revolving around the center of our Galaxy. The Earth is revolving around the Sun and rotating on its axis. All of these motions can add to or subtract from the radial velocity due just to the actual rotation of *Andromeda*. The motions of the Sun and Earth are reasonably well known and are corrected for in the values you have just measured. However, the overall motion of *Andromeda* with respect to us is still in the data and has to be determined. The simplest way to do this is to assume that the center of *Andromeda* is not rotating and that all the motion we see there is systemic, *i.e.*, is due to the overall motion of *Andromeda*.

Record the velocity at the center and enter it in **Table 4**. Alternatively, you can also estimate the systemic velocity by averaging all the velocities you measured, as long as the points are symmetrically placed on either side of the "true" center of the galaxy. Determine the average velocity and enter it in **Table 4.** Indicate in the table whether the systemic velocity implies that *Andromeda* is approaching or receding. (Is the value negative of positive?)

> **Question 2:**   Which velocity do you think gives a better estimate of the systemic velocity of
> *Andromeda*?   The center velocity or the average velocity?   Explain.

Step 4:   Correct for systemic motion

To correct the line of sight velocities for the systemic motion of *Andromeda*, you must simply subtract the systemic velocity from the v_{los} values you previously recorded in **Table 4**.   I.e.,

$$v_{los}^{corr} = v_{los} - v_{systemic} \qquad (6)$$

Do this, using the value you think best represents the systemic velocity, and record the results in **Table 3** in the column labeled v_{los}^{corr}.   Be careful when subtracting negative numbers!

Step 5:   Correct for the inclination of Andromeda

Finally, you have to correct for the fact that the galaxy is tilted and, thus, the measured radial velocities are not the true rotation speeds.   In Part A, you measured the inclination of *Andromeda* and saw how inclination affects the observed line-of-sight velocities.   Now you will apply Eq. 3 to your measure values of v_{los}^{corr} to derive the true rotation speeds of objects it the *Andromeda* galaxy.   Perform this correction on your data, and enter the results in the last column of **Table 3.**

Step 6:   Draw the rotation curve

From your data in **Table 3**, make a plot of the true rotational speed v_{rot} vs. the distance **a** from the galactic center.   Use the graph provided at the end of the lab or, alternatively, use Excel to make the plot.   Draw a smooth curve that best fits all of the points.   This is called the **"Rotation Curve"** of the *Andromeda* galaxy and describes how the orbital speeds of objects in *Andromeda* vary with distance from the center of the galaxy.   Your rotation curve should be reasonably symmetric.   That is, the velocities on one side of the galactic center should be about the same as the velocities of on the other side of the center.   *If you see a*

large systematic difference between the two, then you may have made an error in determining the systemic velocity. Check this!

If you are using Excel, be sure to match the axes and labels to those of the graph provided with the lab. Plot the data points <u>without</u> connecting lines or curves. Then use the Trendline to fit a smooth curve through the data points. Try a polynomial and vary the order until the curve looks something like what you would draw by hand.

C. THE MASS OF ANDROMEDA

You are now ready to transform your rotation measurements into a map of the distribution of mass within *Andromeda*. In the Introduction to the lab, you saw that *Kepler's Third Law* gives the crucial connection between the orbital properties and mass – you will now apply that equation.

Step 1: Prepare the data

The first two columns of **Table 5** already contain the orbital sizes at which you measured the rotation speed, in the original units of kpc and converted to Astronomical Units (AU). From the smooth curve you drew through the data on your Rotation Curve graph (or fit using Trendline), carefully read the values of v_{rot} at these positions and enter them in the third column of **Table 5**. (Alternatively, you can average your measurements for corresponding distances on either side of Andromeda's center.)

Now you'll determine the orbital period P corresponding to each orbital radius. The orbital period is simply the distance traveled in one orbit divided by the orbital speed. The distance traveled is the circumference of the orbit, given by $2\pi a$. Combining these facts, and taking care to consider the units with which the measurements are made, yields:

$$P = 29.7 \frac{a}{v_{rot}} \ years \qquad\qquad (7)$$

where the resultant period will be in units of years if the orbital size a is measured in AU and the orbital speed v_{rot} in km/sec. Using Eq. 7, compute the value of P for each of the positions in **Table 5** and enter them in the fourth column of the table.

Step 2: Compute the "enclosed masses"

Finally, compute the mass enclosed within each orbit. With the proper choice of units, the constant K in *Kepler's Third Law* (see Eq. 1) will be 1. Choosing those units yields:

$$M = \frac{a^3}{P^2} M_\odot \qquad\qquad (8)$$

where the resultant mass will be in units of the Sun's mass if the orbital size a is measured in AU and the period P in years.

Use Eq. 8 to calculate the mass of *Andromeda* interior to each orbit and enter this mass in the last column of **Table 5**. Make a plot of the enclosed mass as a function of distance from Andromeda's center, using the enclosed graph. <u>If you use Excel</u>, match your axes and labels to those of the graph provided with the lab

D. WHAT IT MEANS

Examine the masses that you have calculated.   As you go further out from the center, you include more of the galaxy and the enclosed mass naturally increases.   Eventually you would expect to get to the "edge" of the galaxy, where you have included all the mass.   At this point, the enclosed masses should "flatten out" to a constant value (the total mass of the galaxy), since increasing the distance from the center no longer includes any new material.

> ***Question 4:***   *Careful study of Andromeda has shown that more than 90% of its stars lie in the disk within a distance of ~20 kpc from the galactic center.   Thus, the disk can be said to have a radius of approximately 20 kpc and we would expect the mass distribution to be flattening out by this point.   From your plot, what is the mass within a radius the size of the stellar disk (i.e., 20 kpc)?   What is the mass within the maximum radius you measured?   Do you observe that the mass distribution has flattened out at the edge of the stellar disk?   Have your observations reached the "edge" of Andromeda's mass distribution?   Explain.   If so, how far from the center is the edge?*

You should see that something strange is going on!   Your measurements of the enclosed mass of Andromeda keep on increasing, far beyond the edge of the stellar disk.   This tells you that there is a significant amount of material beyond the stellar disk, even though the galaxy looks dark out there.   Thus, Andromeda is much bigger and more massive than its obvious stellar disk.   The nature of this material, which is called "**Dark Matter**" because it does not emit or absorb light, is unknown and is one of the biggest mysteries in modern astronomy! Evidence for **Dark Matter** has been found in almost all galaxies, including the Milky Way, and is the dominant mass component of galaxies.

> ***Question 5:***   <u>*Based only on the observations you have made*</u>*; do you know the size of Andromeda or its total mass?   Explain.*

> ***Question 6:***   *It has been said that looking at a galaxy in visible light is like looking at the "tip of an iceberg."   Explain this statement considering what you have learned above.*

Table 1: Inclination and Line of Sight Velocity

X	Y	Line of Sight Velocity (v_{los})	Rotational Velocity (v_{rot})	ABS$(v_{los}) \div \cos(i)$
-450	0			
-350	0			
-250	0			
-150	0			
-50	0			
50	0			
150	0			
250	0			
350	0			
450	0			
		Inclination (i) =		

Table 2: Inclination of the Andromeda Galaxy

Major axis (cm)	
Minor axis (cm)	
Inclination i (radians)	
Inclination i (°)	

Table 3: Andromeda Rotation Data

a (kpc)		v_{los} (km/s)	v_{los}^{corr} (km/s)	v_{rot} (km/s)
27.0	NE			
24.0	NE			
21.0	NE			
18.0	NE			
15.0	NE			
12.0	NE			
9.00	NE			
6.00	NE			
3.00	NE			
0.00				
3.00	SW			
6.00	SW			
9.00	SW			
12.0	SW			
15.0	SW			
18.0	SW			
21.0	SW			
24.0	SW			
27.0	SW			

Table 4: Systemic Motion of Andromeda

Velocity at Center (km/s)	
Average Velocity (km/s)	
Approaching or Receding?	

Table 5: Kepler's Third Law Results

a (kpc)	a (AU)	v_{rot} (km/s)	P (yr)	M ($M_\odot$)
3	6.19E+08			
6	1.24E+09			
9	1.86E+09			
12	2.48E+09			
15	3.09E+09			
18	3.71E+09			
21	4.33E+09			
24	4.95E+09			
27	5.57E+09			

Rotation Curve for Andromeda
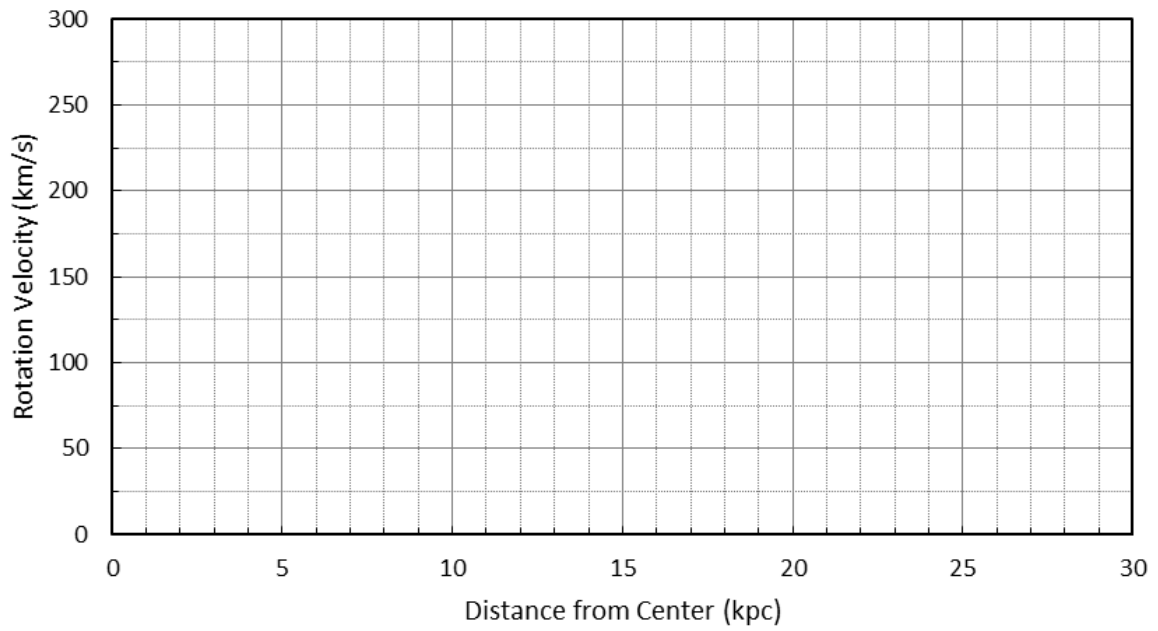
Enclosed Mass for Andromeda
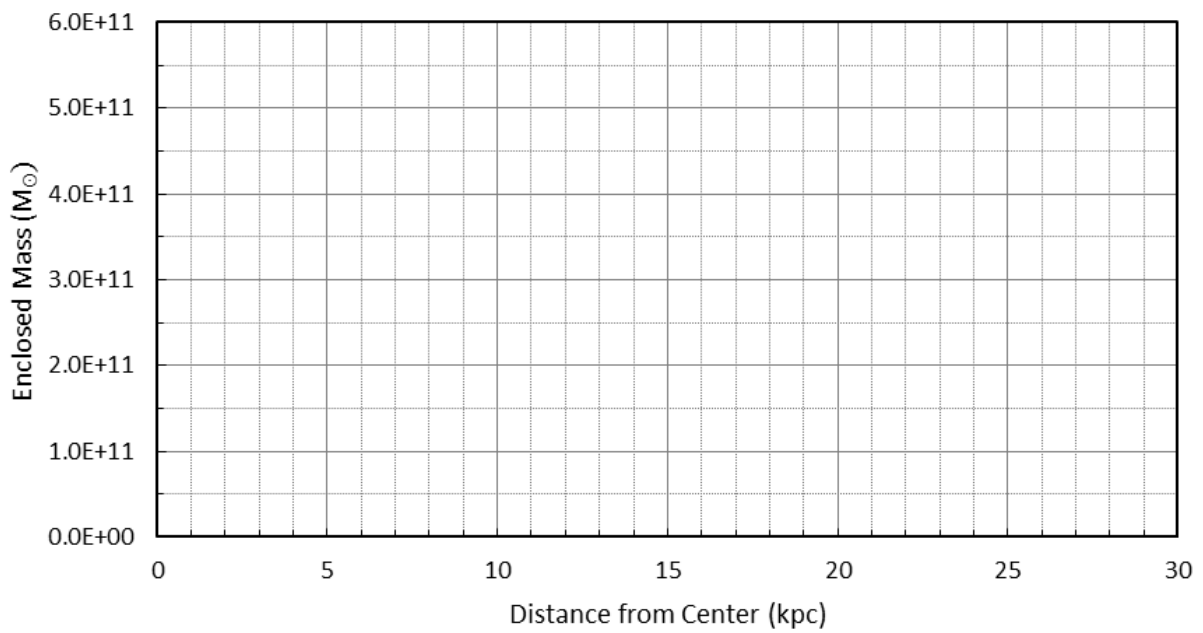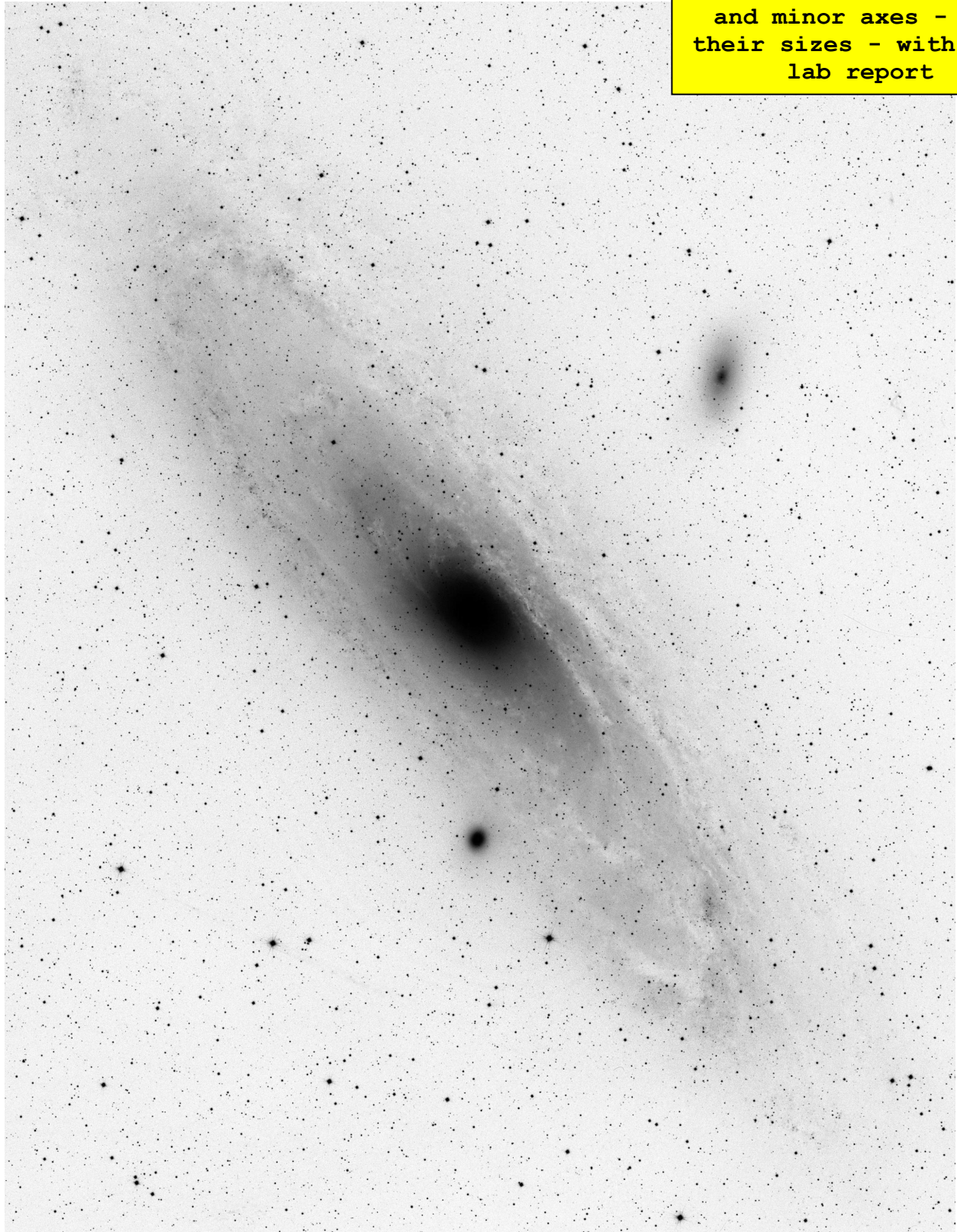
The Andromeda galaxy

Spring 2022
Mendel Science Experience 2151
Astronomy Laboratory - Stars

Lab K
<u>Measuring the Mass of the Andromeda Galaxy</u>

Name: _____

Lab Section: _____

Date: _____

Lab L
The Expansion of the Universe

(Adapted from *Contemporary Laboratory Experiences in Astronomy*)

PURPOSE:

To observe how the apparent recessional velocities of distant galaxies lead to the concept of an expanding Universe, from which the age of the Universe can be estimated.

EQUIPMENT:

Ruler, calculator, Microsoft Excel, and the *CLEA* program ***Hubble Redshift***.
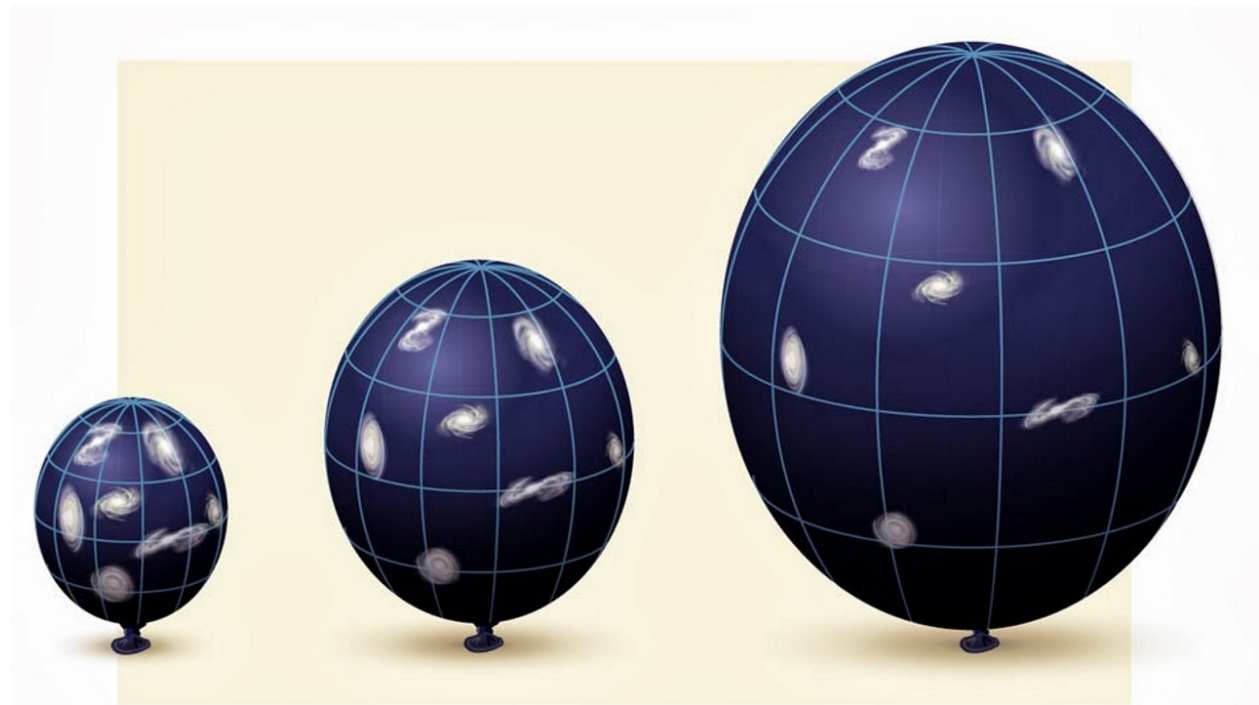
The Expanding Universe
(Credit: Eugenio Bianchi, Carlo Rovelli & Rocky Kolb)

Introduction: The Expansion of the Universe

Early in the1900's, the astronomer V.M. Slipher obtained spectra to study the properties of a number of faint astronomical objects he referred to as "spiral nebulae." He noticed that the spectral lines in many of these nebulae were shifted towards redder wavelengths, as compared with reference lines created in the laboratory. If these "redshifts" are interpreted as a result of the Doppler Effect, then they imply a systematic radial (or line of sight) motion of these galaxies. I.e., Slipher's data implied that most spiral nebulae were moving away from us.

By 1929, Edwin Hubble had established that these spiral nebulae were actually vast galaxies similar to our own Milky Way galaxy. In addition, using techniques he developed for determining the distances to these galaxies, he established that the redshifts of the galaxies were proportional to their distance from the Milky Way. I.e., the more distant the galaxy, the faster it seemed to be moving away from us. **Figure 1**, seen below, is taken from Hubble's 1929 paper and shows a plot of the "recession velocity" versus distance. This type of plot has come to be known as the *Hubble Diagram*. The solid and dashed lines represent Hubble's first estimates of the linear relationship between velocity and distance, now referred to as the *Hubble Law*. This relationship is represented quantitatively as:

$$v = H_0 \times d \qquad\qquad (1)$$

where ***v*** is a galaxy's radial velocity, as inferred from its redshift, and ***d*** is its distance. The parameter ***H₀*** is known as the *Hubble Constant* and represents the slope of the observed relation between ***v*** and ***d*** in the **Hubble Diagram**. It is obtained through observations of numerous galaxies over a range of distances.

If galaxies were randomly distributed throughout the universe and moving in random directions, there would be, on average, just as many of them approaching the Milky Way as receding from it. Yet, Hubble's work convincingly showed that nearly all are actually moving away from our galaxy. What could account for this surprising find? Hubble and others concluded that the Universe itself is expanding in all directions. Galaxies generally appear to be moving further away from one another in much the same way that the galaxy pictures painted on the leftmost balloon in the picture on page 1 of the lab seem to move apart as the balloon is inflated. Once we accept this expansion theory, "running the clock backwards" indicates that, at some time in the very distant past, the components of the universe were very much closer together. If we continue to run time backwards in this manner until everything "touches," then reading our clock will
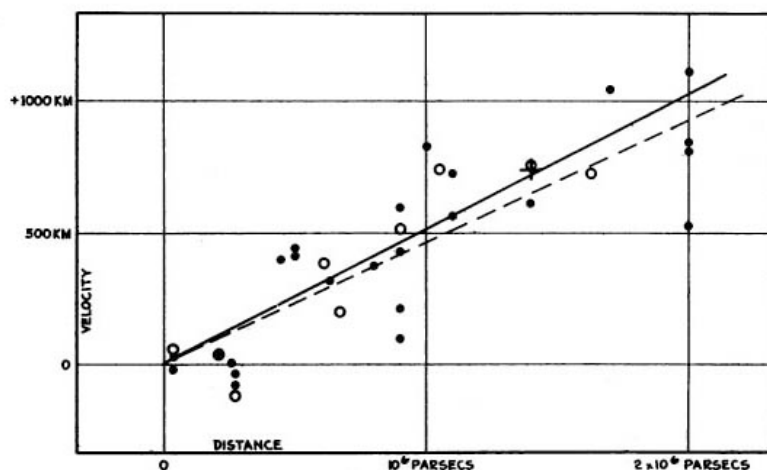

Figure 1: *The original Hubble Diagram*

give us an estimate of the age of the universe.   Today we identify this event as the *Big Bang* – the origin of the Universe and the beginning of time.   The elapsed time since the *Big Bang* depends on the expansion rate of the Universe, which is characterized by the *Hubble Constant*.   As a result, a precise determination of H_0 has been a high priority in astronomy since the 1930's.

Lab Procedure

The software you will use in this lab simulates the operation of a computer-controlled spectrometer attached to a telescope at a large mountain-top observatory.   Using this equipment, you will determine the distances and recession velocities of several galaxies located in selected clusters around the sky.   From these data you will plot a graph of velocity versus distance (as in **Figure 1**) and determine the value of H_0.   Using this value, you will then compute an estimate of the age of the Universe.

So how does the equipment work?   The TV camera attached to your simulated telescope lets you see the galaxies, and you can steer the telescope so that light from a galaxy is focused into the slit of the spectrometer.   You can then turn on the spectrometer and begin to collect photons from the galaxy.   When enough photons are collected, you should be able to see distinct spectral lines from the galaxy (particularly, the *"H"* and *"K"* lines of the element calcium) and you will measure their wavelengths using the computer cursor.   The wavelengths will be longer than the wavelengths of *H* and *K* lines measured in laboratories, because the galaxies are moving away from us and the lines are all redshifted.   The spectrometer also measures the apparent magnitude *V* of the galaxy from the rate at which it receives photons from the galaxy. So, for each galaxy, you will have recorded the wavelengths of the *H* and *K* lines and the *V* magnitude.

A.   COLLECTING THE DATA

Step 1:   Welcome to the Observatory!

Start the ***Hubble Redshift*** program in the "MSE Astrolab" folder   Click **File, Login ...** on the Menu Bar and enter your name.   Click **OK** when finished.   Then click **File, Run** to begin the nights observing.

Step 2:   Prepare the equipment

The screen shows the control panel and viewing window as found in the "warm room" at the observatory. Notice the **Dome** status is "Closed" and **Tracking** status is "OFF".   To begin our evening's work, first open the dome by clicking on the **Dome** button.   When the dome opens you will see a view of the night sky through the finder telescope, showing both stars and galaxies.   The finder scope is mounted on the side of the main telescope and points in the same direction.   It is used to locate the objects we want to measure, because the field of view is much larger than the view in the main instrument.   It is displayed on-screen by a CCD camera attached on the finder scope.   You can switch between the finder scope view and main telescope view by toggling the **Change View** button.

If you look carefully, you will notice that the stars and galaxies seen through the finder scope are drifting slowly towards the west (right) side of the viewing screen.   As you learned in an earlier lab, this is because the Earth is rotating in space.   To be able to point the telescope at any particular object for the time needed to record the desired data, we must counteract this effect.   Click **Tracking** to turn on the telescope's clock drive, which constantly corrects for the Earth's motion.   The tracking status should switch from "OFF" to "ON" and the objects on screen should stop drifting.
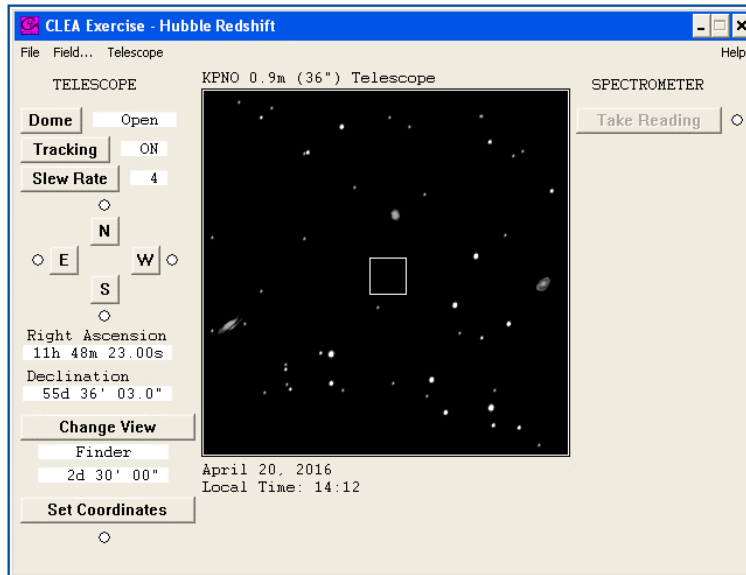
Figure 2: *The simulated telescope control panel*

Step 3:   Select your first target

To find the objects for tonight's observing session, click once on the **Field…** item in the Menu Bar at the top of the control panel.   (Note: This option is *only* available if you are looking through the finder scope.) A small box will open, listing the names of six galaxy clusters that contain the objects we have selected for study tonight, along with the astronomical coordinates of their centers.   These clusters span a wide range of distances and **your goal will be to observe one galaxy from each of five different clusters**.

Click on the cluster name **Ursa Major I**, then click the **OK** button.   The telescope will now "slew" (i.e., move rapidly) to the RA and Dec coordinates of the cluster.   When the telescope arrives at the coordinates, the view window will show a portion of the sky that was electronically captured by the CCD camera attached to the finder scope.   The screen should look like **Figure 2** at the top of this page.   Both stars and galaxies are visible in the view window.   It is easy to recognize bright galaxies in this lab simulation, since the shapes of the brighter galaxies are clearly different from the dot-like images of stars.   But faint galaxies can look like stars, since we can't see their shape. **Be careful! Make sure you observe galaxies and not stars!   If in doubt, ask your instructor!**

The square box in the center of the finder field shows the field of view of the spectrometer. Now move the telescope to place one of the galaxies the within the square box.   Do this by slewing the telescope with the mouse and the **N**, **S**, **E**, and **W** buttons.   Place the arrow on the **N** button and press and hold the left mouse button.   Notice the red light comes on to indicate the telescope is moving in that direction.   As in real observatories, it takes a bit of practice to move the telescope to an object.   You can adjust the telescope's speed by using the mouse to press the **Slew Rate** button.   The slowest rate is 1 and the fastest is 16.

Once the galaxy is within the box, switch to the spectrometer by pressing **Change View**.   Now fine-tune your pointing by moving the spectrometer's "slit" (two short parallel lines) so that it is positioned directly over the galaxy.   Only light passing through the slit will enter the spectrometer.   The more light you get into your spectrometer, the stronger the signal it will detect, and the shorter will be the time required to get a usable spectrum.   Try to position the slit on the brightest portion of the galaxy to maximize the light entering the spectrometer.

Step 4:   Record your first spectrum

When you are satisfied with the position of the galaxy, press the **Take Reading** button.   A new "Spectrometer" window will appear which contains a plot of intensity vs. wavelength (angstroms).   The plot will be empty initially.   To begin taking a spectrum, click **Start/Resume Count**.   The computer will plot the spectrum with the available data as it is accumulated.   Watch the **Signal/Noise** value at the bottom of the window.   We desire this value to be high, which means we have collected more photons from the Galaxy (the signal) than from the sky (the noise).   Click **Stop Count** when the Signal/Noise is *greater* than 20 – the higher, the better.   If you stop the counting before a high-quality spectrum is achieved, you can always restart by clicking **Start/Resume Count** again.

When you are happy with the quality of your spectrum, press **Stop Count**.   At this point, the spectrometer window should look something like **Figure 3** below.   In **Table 1**, or on an Excel spreadsheet supplied by your instructor, record the following information for this observation:

1. Object Name
2. Integration Time (Seconds)
3. Signal/Noise,
4. Apparent Magnitude (*V*)

The appropriate values are listed on the Spectrometer window below the spectrum.
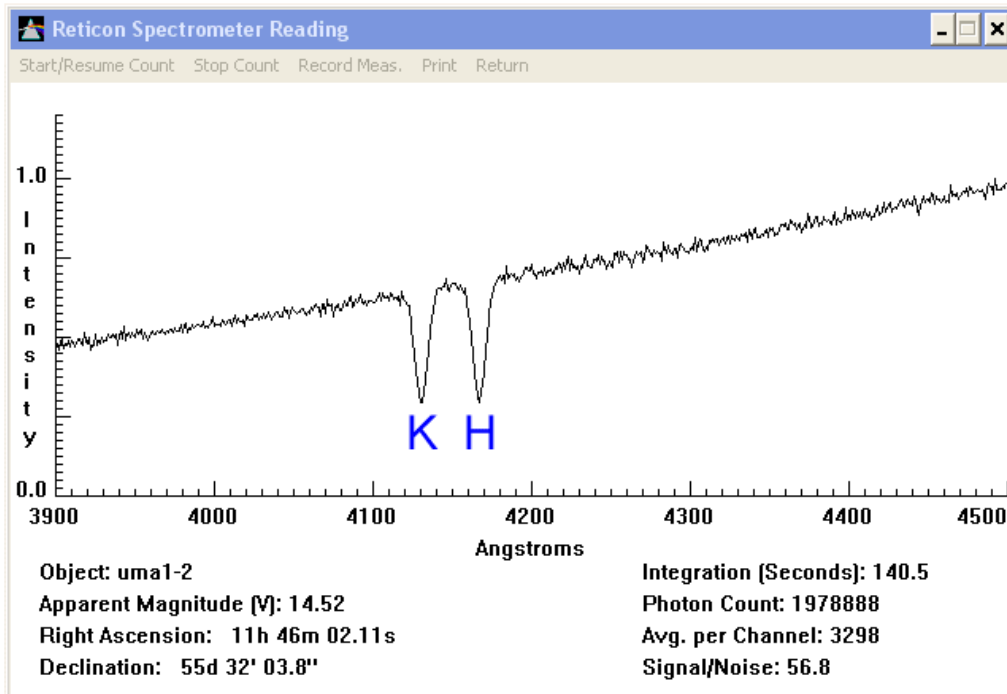
Figure 2: *Your first spectrum!*

Step 5:   Measure the H & K line wavelengths

Look at your spectrum.   The most prominent features in the galaxy's spectrum should be the *H & K* absorption lines of ionized calcium arising from many cool stars within the galaxy.   These lines are approximately 40 angstroms apart, with the *K* line having the shorter wavelength, and they should stand out clearly from the noise.   If not, continue to count photons.   If you are not sure about the data, check with your lab instructor to help you interpret your results.

You now want to measure the wavelength of the centers of the *H & K* lines.   Press and hold the left mouse button, then move the cursor to your estimate of the center wavelength of the *K* line.   Release the mouse button. A tall crosshair will appear at the last mouse position and the wavelength and intensity of the point will appear at the top of the graph.   If you are not satisfied that you have selected the exact center of the *K* line, adjust the cursor position.   When you are satisfied, record the *K* line wavelength in **Table 1**.   Now repeat the procedure for the *H* line.   Note that the line intensities are not important for this lab and do not need to be recorded

Step 6:   Observe the rest of your targets

Now repeat steps 3, 4, and 5 for four more galaxies, one in each of four different galaxy clusters.   Record all your measurements in **Table 1**.

Remember!
Your final set of measurements should consist of one galaxy from each of five galaxy clusters.

B.   COMPUTING VELOCITIES AND DISTANCES

Step 1:   Compute the galaxy redshifts

Compute the *H & K* line redshifts, for each of your target objects and enter the results in **Table 2**.   The redshifts are defined as:

$$redshift = \lambda_{measured} - \lambda_{lab} \tag{2}$$

where the $\lambda_{measured}$ values are from **Table 1.**   The laboratory wavelengths λ_{lab} for the *H & K* lines are 3968.847 Å and 3933.67 Å, respectively.   (Note: your redshift values should all be positive.   If they were negative, they would be called blueshifts!)

Step 2:   Compute the radial velocities

Use the *Doppler Effect* formula to compute the radial velocities *v* implied by the redshifts of the *H & K* lines and enter these values in **Table 2**.   The Doppler formula is given by:

$$v = c \times \left(\frac{\text{redshift}}{\lambda_{lab}} \right) \tag{3}$$

where ***c*** is the speed of light (300,000 km/s).   The resultant velocity *v* has the same units as ***c***.   After all the velocities are computed, average the *H & K* line results for each galaxy and enter that value in the last column of **Table 2**. These mean values will represent the radial velocities of the galaxies.

Step 3: Compute the galaxy distances

You will recall from an earlier lab that the *distance modulus DM* of an object is the difference between its apparent magnitude **V** and its absolute magnitude **M_V**.   *I.e.:*

$$DM = distance\ modulus = (V - M_V) \tag{4}$$

Asin the earlier lab, the *distance modulus* can be converted to distance via the relation:

$$d = 10^{\wedge((DM+5)/5)}\ pc \tag{5}$$

Using **Equation 5**, compute the distances to your five galaxies using your measured **V** magnitudes and the assumed values of **$M_V = -22.0$.**   Enter the results in **Table 3** in units of both parsecs (pc) and megaparsecs (Mpc).

C.   THE HUBBLE CONSTANT AND THE AGE OF THE UNIVERSE

Step 1: Construct a Hubble Diagram and determine H_0

Now plot a *Hubble Diagram* by graphing the average velocity of each of your galaxies (in km/s) on the y-axis vs. their distance (in Mpc) on the x-axis.   Use the graph provided on the last page of the lab manual or, as per your instructor, an *Excel* spreadsheet. If you use the included graph, draw a straight line which passes through the origin [i.e., $(x, y) = (0, 0)$] and best fits all the data points. **The slope of the line is the Hubble Constant (H_0), as defined in Equation 1, with units of (km/sec)/Mpc.**   To calculate the slope of the line, determine a value of **d** and **v** from a point near the *upper right* end of the line you drew (do _not_ use one of the points that you plotted).   The slope of the line will simply be equal to **v/d**, since the line passes through the origin.

If you are using *Excel,* make sure that the axes and labels on your plot match those of the graph included with the lab.   To determine **H_0**, fit a linear Trendline to your data.   In the Trendline options box, check the "Set Intercept" box and set the value to 0.0.   This will constrain the Trendline to pass through the origin. Be sure to have *Excel* print the equation of the Trendline on your graph. The slope of the Trendline is your estimate of **H_0**

Enter your value of the Hubble Constant **H_0**, the current expansion rate of the Universe, in the top row of **Table 4.**   Notice the dimensions of **H_0** (i.e., km/sec/Mpc). **H_0** is represented this way because we typically measure galaxy velocities in units of km/sec and distances in Mpc.   If we had, instead, measured the galactic distances in units of km, then the units of **H_0** would have been km/sec/km which simplifies to 1/sec.   This shows that the natural units for the *Hubble Constant* are inverse time.   Let's now express **H_0** in these more natural units. You could do this by re-plotting your *Hubble Diagram* with galactic distances expressed in km units rather than Mpc, and then re-determining the slope of the best-fit linear relation.   But a quicker way is to just divide your estimate of **H_0** by the number of kilometers in 1 Mpc.   Do this and then enter your result in the second row of **Table 4.**

Step 2: Compute the Age of the Universe

As noted in the lab introduction, the *Hubble Constant* can be used in a very simple way to estimate the age of the Universe.   To see how this is possible, let's start with an analogy.

Suppose you had a driving trip to make and wanted to know how long it would take to complete. To find this time, you would simply divide the distance to be traveled by the speed of the car, using the equation:

$$\frac{distance}{velocity} = time \tag{6}$$

For example, if you must go 120 miles and will be traveling at 60 miles/hour, you come up with a trip time of 2 hours.   Now consider the *Hubble Law* from **Equation 1**, but in a slightly rearranged form:

$$\frac{d}{v} = \frac{1}{H_0} \tag{7}$$

From comparison with **Equation 6**, you can see that the quantity *1/H₀* (usually written as *"H₀⁻¹")* is analogous to the trip time.   What trip?   The trip that the galaxies have taken to reach their current distances, beginning at a time when all the matter in the universe was squashed together!   I.e., the current Age of the Universe.   Thus, we have:

$$The\ Age\ of\ the\ Universe \approx {H_0}^{-1} \tag{7}$$

Using the value of *H₀* you computed above (second line in **Table 4)** compute your estimate of the Age of the Universe in seconds and enter it in the third row of **Table 4**.   Convert the Age into units of years and enter in the last row or **Table 4**.   **Be careful with significant figures!**

Equation 7 contains the "approximately equal to" sign ("≈") because it depends on one key assumption; namely, that the expansion rate of the Universe has always been constant.   Likewise, our calculation of the time needed to drive 120 miles assumed that the speed would always be exactly 60 miles/hour.   It turns out that we now know that the expansion rate of the Universe <u>has</u> varied in the past, but the simple estimate of the Age you just made should be very close to the currently accepted value.

Summary Question: How well did you do?   Go to a web browser and look up the best current estimate of the age of the Universe.   E.g., type "Planck satellite age of universe" in your favorite search engine.   It should return the results from the recent Planck Satellite mission.   What is the currently accepted value?   What is your percent error?

SOME USEFUL QUANTITIES

Speed of light (c)	300,000 km/sec
1 year	31,556,952 sec
1 light year	9.46 x 10^{12} km
1 parsec (pc)	3.26 light years
1 megaparsec (Mpc)	1 x 10^6 pc
1 megaparsec (Mpc)	3.086 x 10^{19} km

Table 1:   Observational Data

| Object Number	Name	Integration time (sec)	Signal to Noise	Apparent Magnitude V	K-line Wavelength (A)	H-Line Wavelength (A)
1						
2						
3						
4						
5						

Table 2:   Galaxy Velocities

Object Number	Name	K-line Redshift (A)	H-Line Redshift (A)	K-Line Velocity (km/s)	H-Line Velocity (km/s)	Average Velocity (km/s)
1						
2						
3						
4						
5						

Table 3:   Galaxy Distances

Object Number	Name	Absolute Magnitude M_V	Distance Modulus	Distance (pc)	Distance (Mpc)
1					
2					
3					
4					
5					

Table 4:   Our Universe

Hubble Constant H_0	km/sec/Mpc
Hubble Constant H_0	sec^{-1}
Age of the Universe	sec
Age of the Universe	years
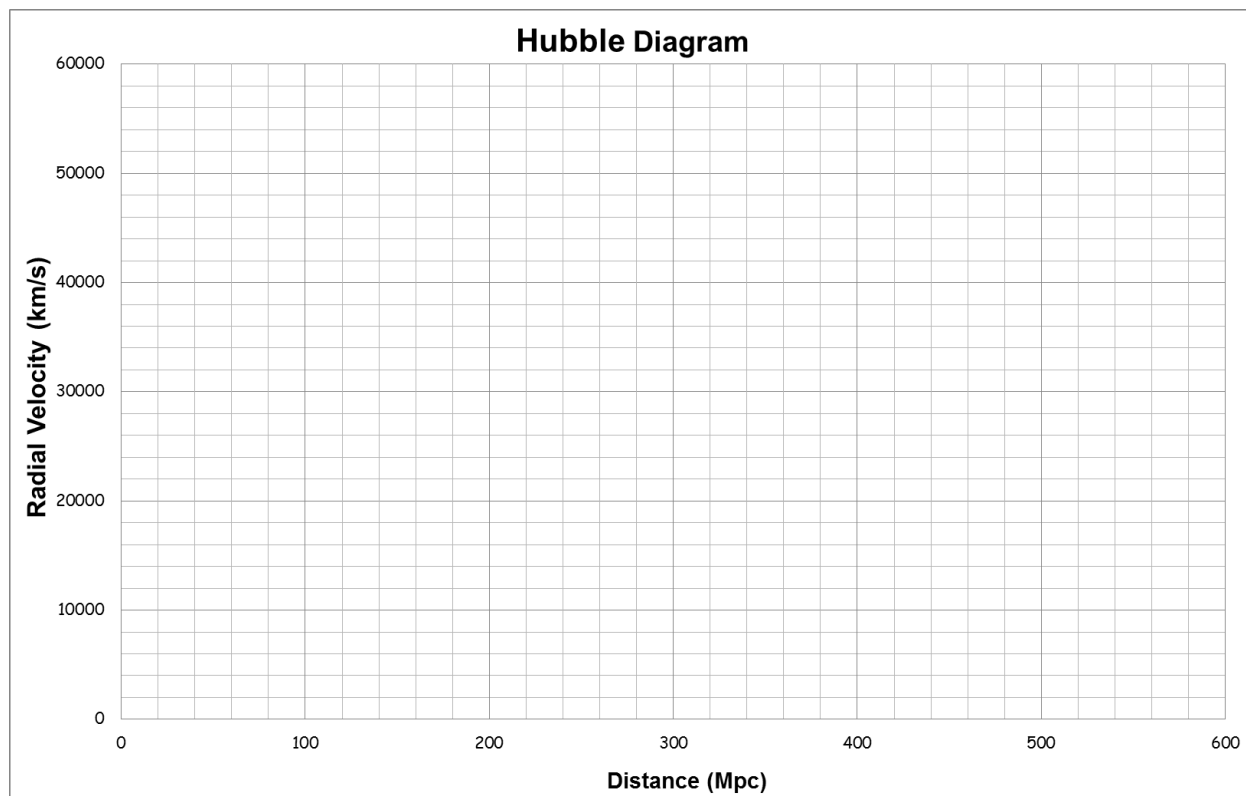
Hubble Diagram

Lab Report Cover Page

Spring 2022
Mendel Science Experience 2151
Astronomy Laboratory - Stars

Lab L
<u>The Expansion of the Universe</u>

Name: _____

Lab Section: _____

Date: _____

Appendix 1

Microsoft Excel Tutorial

Microsoft Excel Tutorial
Creating a Linear Graph with Regression

This tutorial is designed to teach you how to create a graph and perform a linear regression using <u>Microsoft Excel.</u> Excel is a spreadsheet program. This is a program that is designed to take data, usually numerical, and perform various operations on it and output those results as numbers or graphs. While designed primarily for business applications, Excel can also be used for scientific work. The version shown in the pictures will be for Microsoft Office Excel 2013 for Windows.

Launch Microsoft Excel 2013 by selecting its icon.

To begin the process of creating a graph, simply click the mouse in a cell and type the data into that cell. The data can be numbers, words, mathematical formulas, etc. When done, click the mouse in another box and enter the data there. In the example below, we've created two columns: Height and Weight (Height in column B, Weight in column C). This data is from *Lab A: Working with Numbers, Graphs, and the Computer*. When entering data, keep the first couple of rows available for column headings.

When you've entered all of your data, you can create a graph from it. To select the data that you want to graph, click and hold the left mouse button on the upper left cell, then drag the mouse to the lower right cell while holding the button down. The data will appear highlighted as seen below:
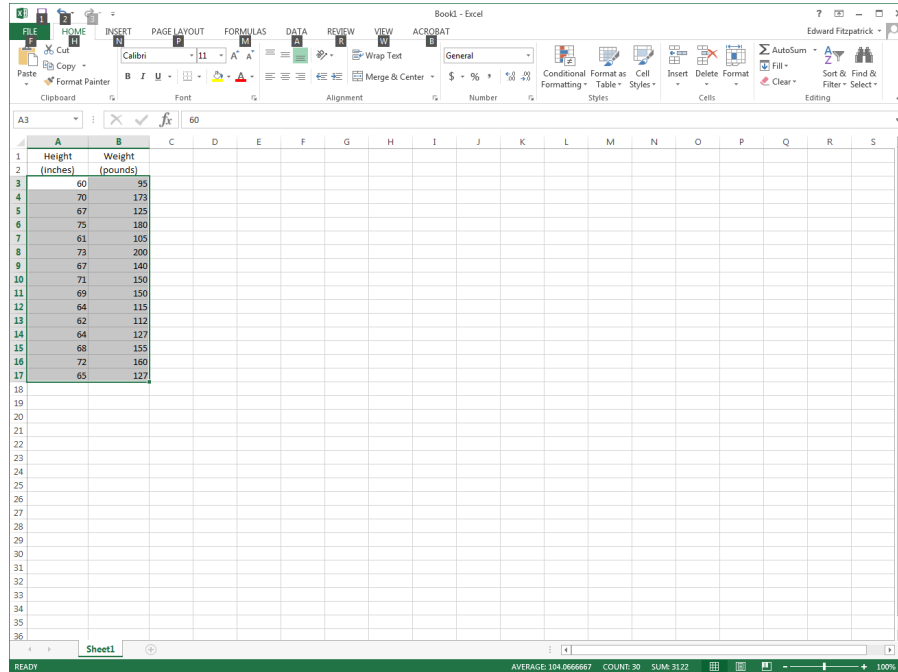
To add a graph to the spreadsheet, you can go to the Insert menu and then choose Scatter in the "Charts" section. For the style of plot, click on the "Scatter with only Markers" type.

This is the usual style of graph that we will use in this class.


Excel -3
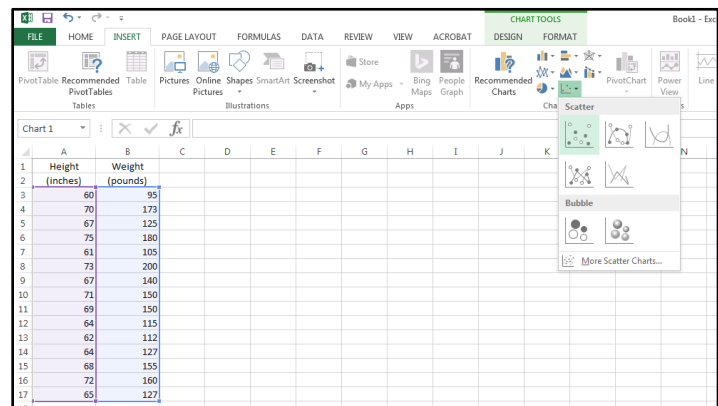
Once your chart appears on the screen, you may want to make changes in how it is displayed or make corrections to your data points. Here, it is necessary to set the range of both the x and y axis to see the data spread out on the page. To do this, click on the graph to highlight it and then select the large plus sign ("+") to the right of the chart:
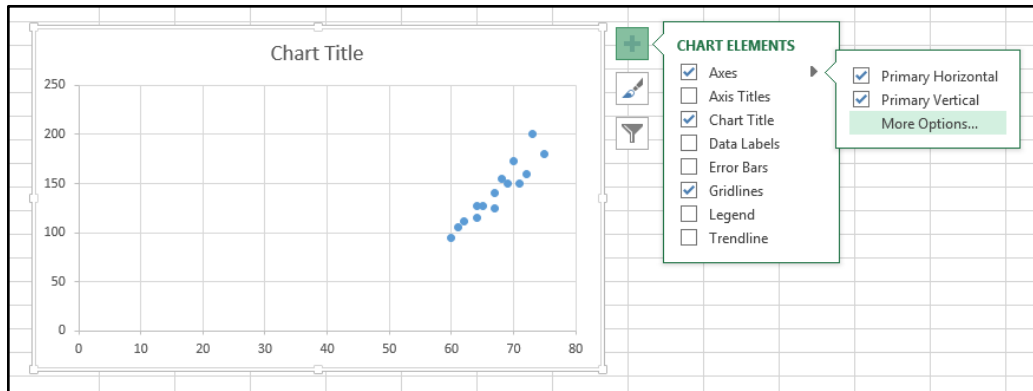
This brings up a number of options which allow you to change the appearance of the chart. To change the range of the graph, click the small arrow next to "Axes" in the "Chart Elements" box and then click "More Options". This brings up the "Format Axis" menu on the right side of the screen:

There are many options available to you (which you will have to explore!) To change the range displayed in the graph, highlight the axis you which to modify (i.e., click on it in the chart), then change the values listed in the "Bounds" section of the menu. You can also adjust the "Units" inputs to change the intervals between the tick marks displayed in the chart. For your Height vs. Weight chart, try x-axis bounds of 55 and 80, with "Major" units set to 5. Try y-axis bounds of 70 and 210, with "Major" units set to 10. Experiment with these values. The goal is for your data to fill the plotting area as much as possible, without wasting space (as in the default chart shown above, where the chart is blank for x-values less than 60 and y-values less than about 90.

It is now time to add axes labels and a chart title. This is done via the "Chart Elements" box. Simply check the "Axis Titles" and "Chart Title" boxes, and text boxes will appear on your chart in which you can type the desire titles and labels. There are options for "prettifying" the titles, which can be accessed through the menus which will appear on the right side of the Excel window. After typing in labels and a title, your chart should look something like:
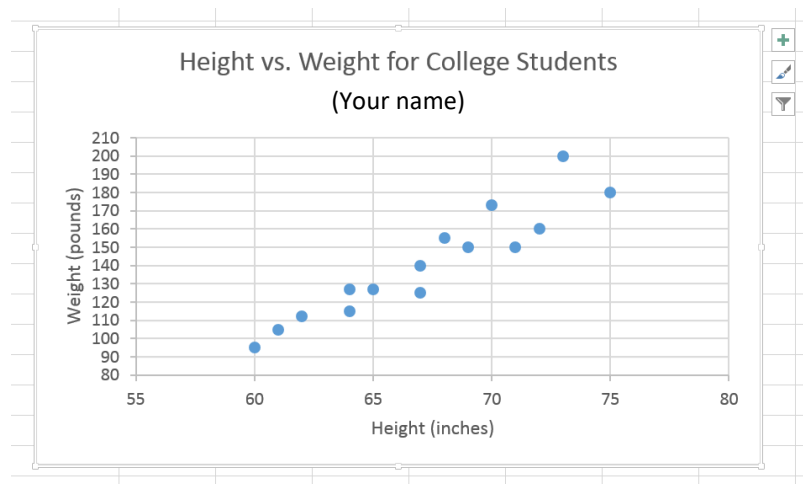
Make sure you enter a descriptive title. It is IMPORTANT to include your name so you can identify your plot at the printer. The axis labels should describe the quantity being displayed (e.g., "Weight") and the units used.

==NOTE:== After creating a beautiful chart, you may discover that you have entered some of the data incorrectly in the spreadsheet. No worries. Simply change the incorrect value in the spreadsheet and the new results will appear automatically on the chart!

TRENDLINES

It is clear from viewing the data in the chart above that height and weight for college students appear to be, on average, linearly-related to each other. I.e., as one increases, the other increases in the same proportion. This relationship can be quantified in Excel by fitting a "Trendline" to the data.

To fit a Trendline to your data, simply check the "Trendline" box in the "Chart Elements" menu. (Remember? Highlight the graph and click the "+" sign to bring up the "Chart Elements" box). Clicking the small arrow next to "Trendline" will bring up the "Format Trendline" menu on the right side of the spreadsheet. There are several options but, in this case, a linear fit seems appropriate. Check the "Linear" option and Excel will perform a linear regression analysis on your data. A line will appear through your data representing the mean relationship between height and weight of college students. Check the "Display Equation on chart" box and the equation of this straight line will appear on the plot. You can drag this equation around to a convenient spot and, also, change the text size and style.

Your final graph will look something like this:
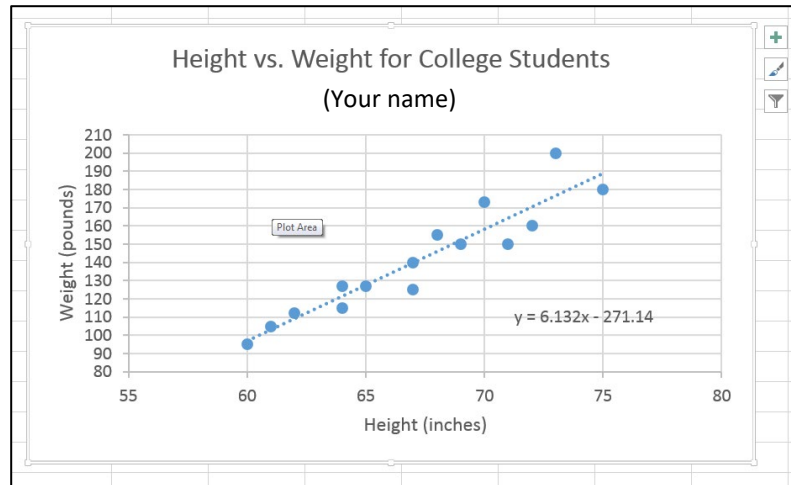
PRINTING

If you need to make a hardcopy of your graph, highlight the graph by clicking on it, then click on the "File" menu in the upper left corner of Excel and chose "Print" to get a view of the whole graph. If it looks OK, then print it and include it with your lab.

NOTE: If you click "Print" in the "File" menu **without** highlighting the graph, the whole spreadsheet will be formatted for printing. Printing this way allows you to include your data tables and charts on the same output page. Elements in the spreadsheet can always be moved around to allow for maximum visibility on the printed output.

Appendix 2

Significant Digits

Significant Digits in a Measurement

Definition:

Significant Digits – All the digits in a numerical measurement that can be justified by the accuracy & precision of the measurement instrument.

✓ **The number of digits reported tells you something about the quality of the measurement:**

- Imagine four length measurements reported like this:

<p style="text-align:center; color:red; font-weight:bold;">
1 meter
1.0 meters
1.00 meters
1.0000 meters
</p>

- These might seem to be the same but actually imply different levels of precision:
  - "1 meter" suggests a "ballpark measurement"
    (could be off by a large fraction of a meter)
  - "1.0000 meters" indicates a very precise result
    (likely to be accurate to nearest millimeter)

✓ **How many digits should be reported:**

- It depends on the properties of the measurement instrument
- The smallest allowable digit in a measurement is the "uncertain digit" or the "estimated digit."
- The "estimated digit" lies between the smallest intervals marked on the measurement device.

✓ **In the sample measurement shown on the next page, either 4.72 or 4.73 are acceptable answers for the length of the pencil. It would <u>not</u> be appropriate to quote a measurement of 4.725.**

Measurement and Significant Digits

4.72 or 4.73 cm
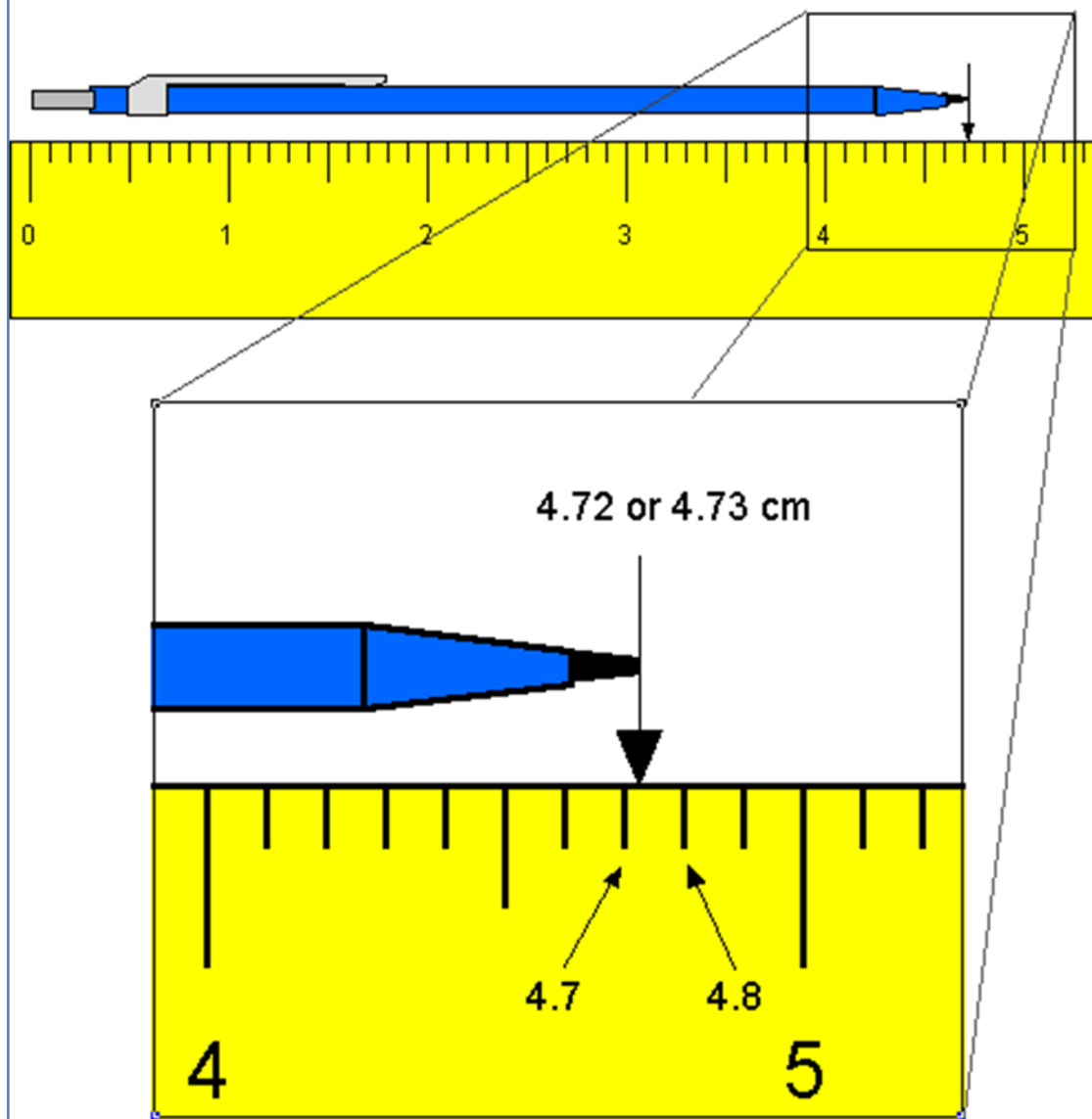
http://chemsite.lsrhs.net/measurement/sig_fig.html

Significant Digits in a Calculation

The Big Idea:

The number of significant digits in the answer to a calculation will depend on the number of significant digits in the original data.

✓ **How to Tell (or Indicate) Whether a Digit is Significant?**

- Non-zero digits are always significant. Thus:

> 22 has two significant digits
> 22.3 has three significant digits

- With zeroes, the situation is more complicated:

  a. Zeroes placed before other digits are not significant; 0.046 has two significant digits.
  b. Zeroes placed between other digits are always significant; 4009 kg has four significant digits.
  c. Zeroes placed after other digits but behind a decimal point are significant; 7.90 has three significant digits.
  d. Zeroes at the end of a number are significant only if they are behind a decimal point as in (c). Otherwise, it is impossible to tell if they are significant. For example, in the number 8200, it is not clear if the zeroes are significant or not. The number of significant digits in 8200 is at least two but could be three or four.

- In scientific notation, all written digits are assumed to be significant. For example:

> 8.200×10^3 has four significant digits
> 8.20×10^3 has three significant digits
> 8.2×10^3 has two significant digits

- In logarithms (or astronomical magnitudes), only the numbers to the right of the decimal point are significant digits.

  For example , consider the number 476.2. It has 4 significant digits. The logarithm of 476.2 should also be expressed using 4 significant digits and would be written as:

$$\log(476.2) = 2.6778$$

✓ **Significant Digits in Multiplication, Division & Trigonometric Functions**

In a calculation involving multiplication, division, and trigonometric functions, the number of significant digits in an answer should equal the least number of significant digits in any one of the numbers being multiplied, divided, etc.

Thus, in evaluating sin(kx), where $k = 0.097$ m^{-1} (two significant digits) and $x = 4.73$ m (three significant digits), the answer should have two significant digits.

Whole numbers have essentially an unlimited number of significant digits. As an example, if a hair dryer uses 1.2 kW of power, then 2 identical hairdryers use 2.4 kW:

1.2 {2 sig. dig.} x 2 {unlimited sig. dig.} = 2.4 {2 sig. dig.}

✓ **Significant Digits in Addition and Subtraction**

When quantities are being added or subtracted, the number of *decimal places* (not significant digits) in the answer should be the same as the *least* number of decimal places in any of the numbers being added or subtracted. For example:

```
      5.67        (two decimal places)
      1.1         (one decimal place)
  + 0.9378        (four decimal place)
      7.7         (one decimal place)
```

✓ **Keep Extra Digits in Intermediate Answers**

When doing multi-step calculations, *keep at least one more significant digit in intermediate results* than needed in your final answer.

For instance, if a final answer requires two significant digits, then carry at least three significant digits in calculations. If you round-off all your intermediate answers to only two digits, you are discarding the information contained in the third digit, and as a result the *second* digit in your final answer might be incorrect. (This phenomenon is known as "round-off error.")

✓ **The Two Greatest Sins Regarding Significant Digits**

1. Writing more digits in an answer (intermediate or final) than justified by the number of digits in the data.

2. Rounding-off, say, to two digits in an intermediate answer, and then writing three digits in the final answer.