Lab 3. Stellar Spectra

1 Introduction

Astronomers have studied stellar spectra systematically for more than 120 years. The earliest extensive work was done visually with eyepiece spectroscopes before photography could be used. In the 1860's Father Angelo Secchi, Director of the Observatory of the Roman College of Jesuits, classified nearly 4000 stars into four groups, based on the appearance of the dark lines in their spectra. The first large-scale, systematic investigation of stellar spectra, however, was not performed until the 1890's when a group of astronomers under the direction of Edward Pickering used photographic spectra to classify over a million stars. (One astronomer, Annie Jump Cannon, herself examined the spectra of over 250,000 objects.) The scheme these astronomers used was alphabetic, A through Q, with A-stars having the strongest hydrogen absorption lines in their spectra and Q-stars having essentially no hydrogen absorption. This original scheme quickly changed, as astronomers began to understand the physics behind the spectral lines: some stellar classes were dropped and others merged, and the letters were rearranged to put the stars in order of *temperature*. Furthermore, astronomers recognized subdivisions within spectral classes. The spectral classification scheme in use today is the MK system, developed in the 1930's and 1940's by Morgan, Keenan, and Kellman. The classification consists of a temperature sequence of the letters O, B, A, F, G, K, and M[1], with subdivisions into tenths, (*i.e.,* A0, A1, A2, ... A9, F0, F1 ...), and a luminosity classification utilizing the Roman numerals V, IV, III, II, and I, with V being used for normal stars like the Sun, and I referring to the extremely luminous supergiant stars. For example, the Sun, a normal star whose spectrum lies two tenths of the way between a G0 star and a K0 star, has a spectral type G2 V.

The basic purpose of spectral classification is to describe the appearance of any stellar spectrum by comparison with *standard stars.* Through the use of detailed mathematical models, astronomers can interpret these spectra and derive quantities such as a star's surface temperature, pressure, surface gravity, electron density, chemical composition, rate of rotation (projected), amount of atmospheric turbulence, magnetic field strength, and even whether the star has active regions such as starspots or flares. However, it is important to remember that although our interpretation of stellar spectra can and does change with time (as our knowledge of physics increases and our models of stellar atmospheres improve), the spectral classification system does not change in any fundamental way, since it is defined strictly by standard stars. For example, a new calibration of the temperatures and luminosities of B3 V stars might be published someday, but a B3 V star will remain a B3 V star.

[1]This sequence has traditionally been memorized as "**O**h, **B**e **A** **F**ine **G**irl, **K**iss **M**e!"

2 Interpreting absorption lines

As mentioned above, one can learn much about a star from its spectrum. As an example, let us consider how an absorption line is affected by the star's temperature. From the previous lab, you learned that an absorption line is produced by a relatively cool gas between an observer and a hot, continuum source. (In the case of stars, this "cool" gas is the star's atmosphere.) Electrons bound within atoms absorb those photons that cause them to jump to a higher energy level, say from level n to level m. Let's call the difference in energy between these two levels, $E_\circ$. Only those photons with a wavelength $\lambda_\circ = hc/E_\circ$ will be absorbed by these atoms; photons not carrying the proper amount of energy will be ignored.

Now what controls how many absorptions by a given element will take place inside a star's atmosphere? Clearly, several factors enter into this calculation. The first is atomic physics — due to the laws of quantum mechanics, some transitions are less probable than others (just because a photon has the proper wavelength to cause an absorption, that doesn't mean it will do so). Next, consider that if an atomic species is to absorb a photon, that species must first exist. For example, if a star is not hot enough to ionize helium, then no matter how probable the transitions of ionized helium are, its absorptions will not be present. Finally, if an electron is to absorb a photon of wavelength $\lambda_\circ$ and go from level n to level m, it must be in level n to start with. If n is the ground level ($n = 1$) of an atom, then chances are there will be plenty of atoms available to do the absorbing. However, if n is large, then the absorptions from that level will be few.

As a specific example, consider the $n = 2$ to $m = 3$ transition of hydrogen. This is called the Hα transition of hydrogen, and corresponds to an wavelength of 6562.8 Å. How will the strength of this line change with temperature? First, we must calculate how many neutral hydrogen atoms are available to perform the absorptions. Hydrogen can be in two states — it can be neutral, or it can be ionized. The energy required to ionize hydrogen is $\chi_i = 13.6$ eV (electron volts), which corresponds to a wavelength of 912 Å, far in the ultraviolet. The more ultraviolet photons shortward of 912 Å a star emits, the more hydrogen atoms it will ionize. Thus, the hotter the star, the smaller the fraction of neutral hydrogen. An expression that relates the fraction of ionized hydrogen to neutral hydrogen $N(H\ II)/N(H\ I)$ is called the Saha equation, and is

$$\frac{N(H\ II)}{N(H\ I)} N_e = \left(\frac{2\pi m_e kT}{h^2}\right)^{3/2} \frac{2\, u_{II}}{u_H(T)} e^{-\chi_i/kT}, \tag{1}$$

where m_e is the mass of the electron, N_e the density of electrons in the star's atmosphere, and u_{II} and $u_H(T)$ arise from the quantum mechanical statistical weights, and are called partition functions. (Note here the astronomer's notation used for expressing ionizations: the roman numeral I after an element indicates that the species is neutral, *i.e.,* it has all its electrons; a II indicates the element is singly ionized; III means doubly ionized, *etc.*) By combining the constants, using the ideal gas law to replace the electron density with pressure, and taking the logarithm (base 10), equation (1) can be simplified to

$$\log \frac{N(H\ II)}{N(H\ I)} = (-5040/T)\,\chi_i - \log P_e + 2.5 \log T - 0.48 + \log \frac{2u_{II}}{u_H(T)}, \tag{2}$$

where χ_i is measured in eV. For ionized hydrogen, $u_{II} = 1$.

Next, we must calculate what fraction of the neutral hydrogen atoms actually have their electron in the $n = 2$ state. Normally, hydrogen atoms prefer to be in their ground ($n = 1$) state, and any hydrogen electron that finds itself in an excited state will decay downward within $\sim 10^{-7}$ seconds. Hence, under normal conditions, you would expect that the fraction of hydrogen atoms in $n = 2$ would be minuscule. However, within a stellar atmosphere, there are frequent collisions between atoms, and in these collisions, electrons can be bumped back up to higher levels. Again, the ability of collisions to populate the upper levels of hydrogen depends on the temperature of the star: the hotter the star, the more violent and frequent the collisions between atoms, and the easier it is for electrons to become excited. In thermodynamical equilibrium, the Boltzmann equation gives the relative populations of each energy level,

$$\frac{N(m)}{N(n)} = (g_m/g_n)\, e^{-\chi_{m,n}/kT}, \tag{3}$$

where g is called the quantum mechanical statistical weight, and $\chi_{m,n}$ is the energy difference between level m and level n. In the case of the Hα line of hydrogen, $\chi_{3,2} = 1.89$ eV, so after some simplification, equation (3) becomes

$$\log \frac{N(n=2)}{N(H\ I)} = (-5040/T)\chi_{3,2} + \log \frac{g_{3,2}}{u_H(T)}, \tag{4}$$

where $u_H(T)$ is again the partition function and $g_{3,2} = 2.25$.

By combining equations (2) and (4), X, the number of hydrogen atoms available to absorb an Hα photon in any star can be calculated. For strong lines, the amount of absorption goes as

$$W = X^{2/5}. \tag{5}$$

where X is the number of atoms doing the absorbing.

3 The lab

In this lab, you are presented with digital spectra taken at Kitt Peak National Observatory. Using these spectra, you will identify some of the elements present in stellar atmospheres and examine how the strength of the Hα absorption changes with spectral type.

Fundamental to this latter measurement is the concept of *equivalent width.* As you will see in the spectra (and is demonstrated in Figure 1), within an absorption line, the light from a star

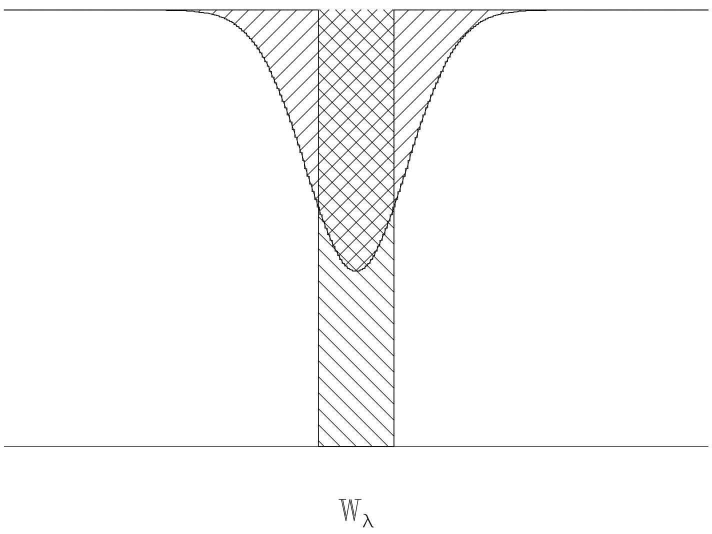

Figure 1: Definition of equivalent width.

is decreased, but does not go to zero. Some lines have strong, wide absorptions, while others absorptions are relatively shallow and narrow. To place all these line shapes on the same scale, astronomers use a measure called equivalent width. *The equivalent width of an absorption line is defined as the width of a rectangle whose base is at zero intensity and whose area is the same as that of the true absorption line.* Equivalent widths are measured by defining the stellar continuum on each side of the absorption, measuring the number of missing counts in the absorption line (an area) and dividing by the average height of the continuum.

In this computer lab you will find a set of spectra for stars of different spectral types. You will plot the spectra, identify some of the absorption lines (and the elements doing the absorbing), and measure the equivalent width of Hα as a function of spectral type.

3.1 Procedure

1. You will need two new IDL programs for this lab, *reader.pro* and *equiv.pro*. Copy them both to your directory by typing

   cp /afs/psu.edu/users/m/a/mab22/lab3/reader.pro   reader.pro

   followed by

cp /afs/psu.edu/users/m/a/mab22/lab3/equiv.pro equiv.pro

2. Enter IDL by typing **idl** (small letters) from your xterm. Remember, most instructions are case sensitive. Also, if you get into trouble in IDL, type the word **retall**.

3. Before you get going any further, you have to load a software library. Type

   astrolib

4. Table 2 lists the names of the files containing the stellar spectra. Each file contains 2799 rows of numbers, each containing a wavelength and a flux. You can view one of these files, say the file containing the A1 V star by typing

   $ page /afs/psu.edu/users/m/a/mab22/lab3/star.a1v

   When you are tired of looking at the numbers, type **q**.

5. Introduce IDL to your measurement programs. First type

   .run reader

   This is a very simple routine that will allow you to read in a 2-column array. As always, you can view this file by typing

   $ page reader.pro

   Next, introduce IDL to the program **equiv** by typing

   .run equiv

   Finally, re-introduce IDL to the program you used last week,

   .run pixel

6. Choose one of the file names listed in Table 2. Read the file by typing

   reader, *filename,* **2799, wave, flux**

   where *filename* is the name of the file you chose, and MUST BE ENCLOSED IN QUOTE MARKS. This command says "read in the 2799 rows of *filename*, put the first column of numbers into **wave** and the second column of numbers into **flux**." (Actually, you can use any variable name at all for **wave** and **flux**. I chose those names since the first column gives wavelength (in angstroms) and the second column gives the observed flux (in ergs/cm^2/s/Å.).

7. Plot the spectrum, either by typing

   plot, flux or **plot, wave, flux**

   The first option will plot the flux values against pixel on the detector; the second instruction will plot flux against wavelength. Remember, you can scale, expand, and label plots using the same options you learned last week. Also, if you only want to plot only a subsection of the graph, you must say

   plot, flux(*x1*:*x2***)** or **plot, wave(***x1*:*x2***), flux(***x1*:*x2***)**

   where *x1* and *x2* are the first and last PIXEL values you wish to plot.

8. Use the line lists provided at the end of this lab to identify the prominent absorption lines in the star. What elements are you seeing? In order to measure accurate wavelengths, you may want to use the **pixel** program from last week. You can type

   pixel, wave, flux

   to display and measure lines over the whole spectrum,

   pixel, wave, flux, $x1$, $x2$

   to display and measure the spectrum between PIXEL values $x1$ and $x2$, or

   pixel, wave, flux, $x1$, $x2$, **'w'**

   to display and measure the spectrum between WAVELENGTH values $x1$ and $x2$. Notice that some of the wavelengths you measure may be off slightly. Why do think this may be?

9. Repeat items 5 through 7 for a selection of stars of different spectral classes. How do the prominent absorption lines vary with spectral type. (For this part of the lab, you need not do every star — maybe one from each spectral class will do.)

10. Now focus in on the Hα line, and measure the equivalent width of Hα in each spectrum provided. You can do this by typing

    equiv, wave, flux

    Again, if you want to zoom in on the spectrum between $x1$ and $x2$, you can type

    equiv, wave, flux, $x1$, $x2$, **'w'**

    where the optional argument **'w'** tells IDL to interpret $x1$ and $x2$ as WAVELENGTHS instead of PIXELS. Plot the equivalent width of Hα as a function of spectral type.

11. Now, focus in on the Hα line. Use **equiv** to measure the equivalent width of Hα in each spectrum provided. Plot these widths as a function of spectral type.

12. Table 1 in this lab gives the partition function U for Hydrogen. Table 2 lists, as a function of spectral type, the effective temperatures and electron pressures in the atmospheres of stars. Use these tables and equations (2), (4), and (5) to predict the spectral type where the maximum amount of absorption should be. How does this compare with the observations?

Table 1: Values of the partition function

$T_{\rm eff}(K)$	$u_H(T)$
<10000	2.00
15000	2.05
18000	2.25
21000	2.84
23000	3.57
25500	5.00
27000	6.10

Table 2: Filenames for Stellar Spectra

Filename	Spectra Type	$T_{\text{eff}}(K)$	$\log_{10} P_e$ (dynes-cm^{-2})
/afs/psu.edu/users/m/a/mab22/lab3/star.o5v	O5 V	54,000	3.3
/afs/psu.edu/users/m/a/mab22/lab3/star.o7v	O7 V	43,300	3.2
/afs/psu.edu/users/m/a/mab22/lab3/star.o9v	O9 V	37,800	3.1
/afs/psu.edu/users/m/a/mab22/lab3/star.b0v	B0 V	29,200	3.0
/afs/psu.edu/users/m/a/mab22/lab3/star.b3v	B3 V	17,600	2.8
/afs/psu.edu/users/m/a/mab22/lab3/star.b6v	B6 V	14,300	2.6
/afs/psu.edu/users/m/a/mab22/lab3/star.b8v	B8 V	12,300	2.5
/afs/psu.edu/users/m/a/mab22/lab3/star.a1v	A1 V	9,330	2.4
/afs/psu.edu/users/m/a/mab22/lab3/star.a3v	A3 V	8,750	2.3
/afs/psu.edu/users/m/a/mab22/lab3/star.a6v	A6 V	8,120	2.1
/afs/psu.edu/users/m/a/mab22/lab3/star.a8v	A8 V	7,930	2.0
/afs/psu.edu/users/m/a/mab22/lab3/star.f0v	F0 V	7,350	1.9
/afs/psu.edu/users/m/a/mab22/lab3/star.f3v	F3 V	6,850	1.6
/afs/psu.edu/users/m/a/mab22/lab3/star.f5v	F5 V	6,700	1.4
/afs/psu.edu/users/m/a/mab22/lab3/star.f7v	F7 V	6,400	1.3
/afs/psu.edu/users/m/a/mab22/lab3/star.f9v	F9 V	6,170	1.1
/afs/psu.edu/users/m/a/mab22/lab3/star.g0v	G0 V	6,050	1.0
/afs/psu.edu/users/m/a/mab22/lab3/star.g2v	G2 V	5,770	0.8
/afs/psu.edu/users/m/a/mab22/lab3/star.g4v	G4 V	5,700	0.7
/afs/psu.edu/users/m/a/mab22/lab3/star.g6v	G6 V	5,590	0.6
/afs/psu.edu/users/m/a/mab22/lab3/star.g9v	G9 V	5,340	0.5
/afs/psu.edu/users/m/a/mab22/lab3/star.k0v	K0 V	5,240	0.5
/afs/psu.edu/users/m/a/mab22/lab3/star.k4v	K4 V	4,600	0.1
/afs/psu.edu/users/m/a/mab22/lab3/star.mov	M0 V	3,750	-0.2
/afs/psu.edu/users/m/a/mab22/lab3/star.m5v	M5 V	3,200	-0.6

Table 3: An incomplete list of astrophysically important spectral features

Identification	Wavelength (Å)	Identification	Wavelength (Å)
He I	10830	O II	4649
C I	10691	N III	4641
Si I	10689	N III	4634
Si I	10627	Mg I	4571
Si I	10603	He II	4541
Si I	10371	Mg II	4481
He II	10124	He I	4472
He II	10120	He I	4471
Na I	9961	He I	4388
C III	9710	Fe I	4384
O I	8446	H I γ	4340
He II	8237	O II	4318
O I	7774	G-band	4300
He I	7065	Ca I	4227
H I α	6563	He II	4200
Na I (D)	5896	H I δ	4102
Na I (D)	5890	He II	4100
He I	5876	N III	4097
He II	5412	Si IV	4089
Mg I (b)	5175	O II	4073
Mg I	5173	He I	4026
He I	4922	H I ϵ	3970
H I β	4861	Ca II (H)	3968
He II	4686	He I	3965
C IV	4658	Ca II (K)	3934
O II	4650	He I	3820
C III	4650	O III	3760