Chapter 14: **Stellar spectra and the Hertzsprung-Russell diagram.**

- stellar spectra; lines; spectral types
- the HR and color-magnitude diagrams
- stellar properties
- sample problems

Stellar Spectra

The interiors of stars can be treated as black bodies. If you were deep inside a star the light coming at you from all directions would look like a black body spectrum characterized by your local temperature. And that black body temperature would be in equilibrium with the temperature you might determine from the general level of ionization of the atoms around you or from their kinetic energies. There would be a slight difference, of course, because looking inward it is hotter than looking outward. In the stellar atmosphere conditions are different: outward is the cold blackness of space, inward the hot interior of the star. Temperatures in stellar photospheres are lower, ranging from ~3,000 to ~30,000 K, densities are lower, and atoms are more likely to be neutral and electrons more likely to be in lower energy levels than is the case in the stellar interior. The temperatures, based on kinetic energies or electron excitation levels or the peak of the spectrum, are no longer all in equilibrium.

Electrons in atoms can absorb particular wavelengths of light, jumping up to higher energy levels. If the energies are high enough, ionization occurs, i.e., the electrons are kicked off entirely. Absorption of specific wavelengths and of the broad range of wavelengths giving rise to ionization are two of the main contributors to *opacity*, i.e., there are colors where the stellar material is relatively opaque as opposed to transparent. Opacity is a function of wavelength; it is a combination of density — there has to be something there to absorb the light — and the propensity of the material that is present to absorb (or scatter or otherwise interfere with) light at that wavelength. The stellar photosphere is the layer from which photons stand a better than even chance of escaping. Some wavelengths, which correspond to particularly likely energy transitions given the mix of atoms present in the photosphere — in other words, wavelengths at which the opacity is high — will continue to experience absorptions for some hundreds of kilometers farther outward than wavelengths that can't be absorbed in the photosphere. The result is a stellar spectrum that closely resembles that of a black body with absorption lines due to those last high-probability energy transitions.

Absorption lines are not all equally strong. Different transitions remove different amounts of energy from the overall spectrum and we need a means of characterizing the strength of a line. Let's look at a chunk of the stellar spectrum, perhaps only a few nanometers wide, across an absorption line and at the shape of the line profile. The following shows a representative absorption line profile. The line has a core and wings, which flatten out and approach the flux level of the nearby continuum. The line may have other, weaker, lines very nearby; in complicated regions of the spectrum it can be difficult to distinguish one line from another or to determine the level of the continuum. Our line here is not too messy.

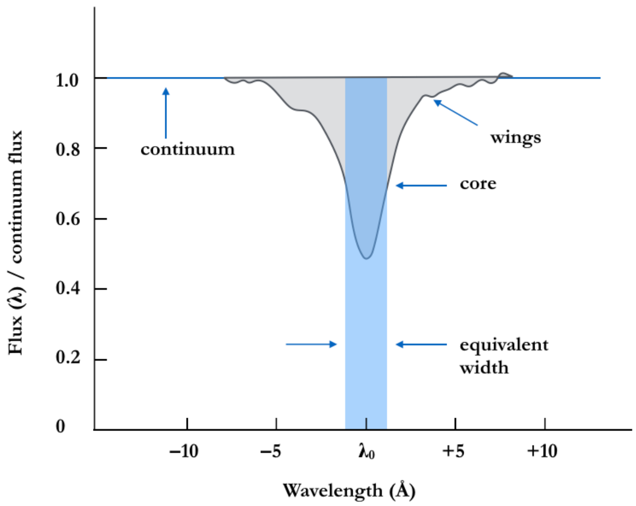

Figure 14.1: Representative absorption line profile.

We are representing the continuum flux level as being flat simply because this is such a small region of the spectrum. In reality the line is superposed on the overall curved spectrum of the underlying approximately black body. Notice that the line core is not totally black, i.e., the flux at the center does not drop to zero. This is principally because energy absorptions are followed by emissions; the energy being absorbed is directional, in the sense that it's only arriving from the interior of the star, while the energy can be reemitted in any direction, including outward toward us. Even emissions from very strong lines, where the photons only escape from the lowest temperature / top of the photosphere, are not coming from a region with zero temperature, and that means they can't have zero flux. The *equivalent width* of the line is the width of a rectangle of continuum height having the same total area as the area between the actual line profile and the continuum. In other words, the equivalent width gives an indication of the fraction of the available flux that is missing due to the absorption line. Describing the strengths of absorption lines usually means describing their equivalent widths.

Every different ionization state of every element has its own specific set of possible electron energy levels. For an absorption line to be present, in addition to having light of the right wavelength, the relevant ionization state must be present and a substantial fraction of those atoms or ions must have their electrons in the correct lower level from which they can absorb the energy that corresponds to the wavelength of that absorption line. You have already read that the hydrogen Balmer series of absorption lines arise from neutral hydrogens with their electrons in the second energy level. Lyman-α, the absorption that will boost the electron from the ground state of hydrogen into that second energy level, is in the ultraviolet. Stars that are too cool won't have enough ultraviolet to boost substantial numbers of hydrogen electrons into the second energy level and thus cool stars won't show strong Balmer lines. Stars that are too hot will have so much ultraviolet that too high a fraction of their hydrogens will be ionized, thus also resulting in very few electrons in the second energy level; hot stars also will not show strong Balmer lines.

The various elements and their dominant ionization stages have electrons that are more or less hard to excite depending on how tightly those electrons are bound by the positively charged nucleus. Relative to hydrogen, the outer elements of heavier elements — metals — such as iron or magnesium are relatively easier to excite because the outermost electrons are shielded from much of the positive charge by inner, more tightly bound, electrons. Cool stars have lots of metal lines. Helium electrons, on the other hand, are hard to excite and helium is quite hard to ionize. Helium and He^+ lines show up in the spectra of hot stars. Here are a few ionization energies, which give an indication of how tightly bound the electrons are.

1st ionization energy	2nd ionization energy	3rd ionization energy	4th ionization energy
H: 13.6 eV			
He: 24.6 eV	54.4 eV		
Mg: 7.6 eV	15.0 eV	80.1 eV	109.3 eV
Si: 8.2 eV	16.3 eV	33.5 eV	45.1 eV
Fe: 7.9 eV	16.2 eV	30.7 eV	54.8 eV

Line profiles are not infinitely skinny but are broadened by several mechanisms. One reason that the lines have width is that energy levels are not precisely defined. This can be related to the Heisenberg Uncertainty Principle. You are probably most likely to have seen this expressed as $\Delta x \Delta p \geq \hbar/2$, meaning that the product of the uncertainty in the value of a particle's position and the uncertainty in the value of its momentum cannot be arbitrarily small but is limited to the value $\hbar/2$, where $\hbar = h / 2\pi$ is called the reduced Planck's constant. This is a statement about the nature of quantum systems, not about our technology; in other words, it's not simply that we are unable to measure precisely enough. Taking momentum as mass times velocity, consider the units on position and momentum: m · (kg·m/s). The joule - second, expressed in fundamental quantities, has units of [(kg·m^2/s^2)· s]. The uncertainty principle may also be expressed as $\Delta E \Delta t \geq \hbar/2$ — the units are the same. The implication for stellar spectra is that electron energy levels are inherently fuzzy. The longer an electron is likely to stay in an energy level, the more tightly constrained that energy level is. For most transitions the electron's lifetime in a particular energy state is very short, on the order of 10^{-8} s, and the energy levels are thus not totally sharp. There is a distinct probability that the electron will absorb an energy that is slightly off of the expected energy, or, in terms of our line profile, at a wavelength that is slightly off of λ_0. This is called natural broadening.

A second set of reasons that lines are broadened has to do with motion. Stars rotate. Unless we are observing a star pole-on, this means that some parts of the star are coming toward us and some are moving away. Some parts of the star will be moving transversely to our line of sight but many of the individual absorptions that go into making up the absorption line are Doppler-shifted. This is called rotational broadening. When a star rotates rapidly, the spectral lines tend to be broad and also shallower than they would be in the absence of rapid rotation. The number of photons that can be absorbed in a line depends in large part on the number of atoms available to absorb that wavelength. You could imagine speeding up the rotation rate of the star, but that won't change the number of absorbers, so the equivalent width of the line doesn't need to increase.

Even if the star were not rotating, though, there would be Doppler broadening on a smaller scale because of the kinetic energies of the atoms doing the absorbing. Some atoms are moving toward us, some cross-ways, and some away meaning that we will see some absorptions slightly blue shifted, many not shifted, and some slightly red shifted relative to the expected absorption at λ_0. Higher temperatures mean faster atoms and thus more broadening. The following sketch is an attempt to illustrate this by drawing the line profile used above, when we were considering equivalent width, as if it were composed of many individual little lines, Doppler shifted away from line center. Line center is still the most likely wavelength, and the little lines are thus shown as larger as they approach line center.

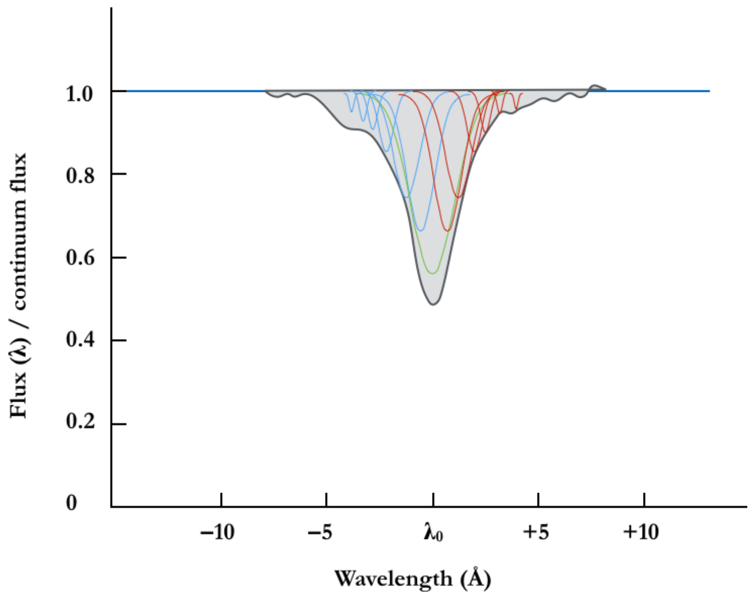

Figure 14.2: Doppler broadening of a representative absorption line profile.

A third reason for line broadening that is a bit more subtle involves the density of the star's atmosphere. The energy levels of an atom that's absorbing will be affected by the electric fields of neighboring atoms. The more nearby atoms, the more effect. The result is that absorption lines in very distended giant stars, where the atmospheric density is low, will all be thinner than the absorption lines in the spectra of stars with denser atmospheres. This is called pressure broadening.

Many absorption lines will also split (not just broaden) in the presence of a magnetic field. Energy levels are actually made up of many close or overlapping sub-levels, some of which may shift slightly to higher or lower energies if the absorbing atom is in a magnetic field. The fact that some lines split is useful for determining the strength of the magnetic field in sunspots; the stronger the field, the more the lines separate in wavelength.

Our confusing ordering of spectral types (OBAFGKM, etc., about which more below) arises from the fact that the first attempts to sort spectra of various stars and describe similarities and differences between stars' spectra were made several decades before anyone knew about electrons and their energy levels. The first significant attempt to organize stellar spectra was made by Fr. Angelo Secchi (Italian; 1818 - 1878). In the 1860s he began sorting spectra into three main classes: Class I were stars with strong hydrogen lines; class II, stars such as the Sun in which hydrogen lines are present but less strong and which show lots of lines of metals; class III, red stars whose spectra show weak hydrogen lines and complex bands of lines (which we would now attribute to molecules). With more spectra, Secchi added class IV, stars with stronger than average lines of carbon, and, finally, class V, stars whose spectra show emission lines.

In the 1880s Edward Pickering and his staff astronomers at the Harvard College Observatory began work on a survey of stellar spectra, producing the Henry Draper Catalogue; many stars today are still identified by their HD numbers, based on that catalogue. Williamina Fleming classified the majority of the spectra for the first Draper catalog, using letters (A, B, etc.) rather than Secchi's Roman numeral scheme. Annie Jump Cannon (1863 - 1941) studied physics and astronomy at Wellesley College and also for a time at Radcliffe. Pickering hired her in 1896 and she soon began work on the Draper catalogue, for which Pickering hoped to classify every star with a photographic (i.e., blue-sensitive) magnitude of 9 or brighter. Cannon sorted spectra based on the strengths of the Balmer lines but organized so that spectra of all elements varied smoothly from one type to the next.

Many talented women worked at the Harvard College Observatory around the turn of the twentieth century. Women's access to a college education was on the rise in the U.S., but access to professional astronomical positions could be hard to come by. A position at the Harvard College Observatory gave women a chance to do astronomy, but mostly doing tedious work for low wages with very little hope of advancement. Men, not women, were permitted to travel to distant telescopes to take the data the women back in Cambridge were analyzing. In her lifetime Cannon classified approximately half a million stellar spectra with amazing consistency. She did receive some recognition for her work, for instance becoming the first woman to receive an honorary doctorate from Oxford University.

Only **B**umbling **A**stronomers **F**orget **G**enerally **K**nown **M**nemonics **L**ike **T**his. The main spectral classes are labelled OBAFGKM. With the work of Cecilia Payne (later Payne-Gaposchkin; 1900 - 1979) in the early 1920s it became clear that this ordering of stellar spectral classes is fundamentally a temperature sequence. Payne was British and came to the U.S. after college to do graduate study at Radcliffe (the women's college associated with Harvard) where she produced what has justifiably been called one of the most brilliant astronomy theses ever written. Payne was able to combine work by Indian physicist Meghnad Saha on the likelihood of ionization of the elements with the observations of stellar spectra to show that stars are overwhelmingly composed of hydrogen even if, as in the hot O stars or cool K and M stars, the Balmer lines are not very strong. The spectral classes are further subdivided into 10 subtypes. For example, the Sun is a spectral type G2; cooler than the Sun we have subtypes G3, G4, G5. . .G9, K0, K1, etc. The hotter stars, types O-B-A, are often called early-type stars and the cooler stars are called later-type. The classes BAFGKM mostly align with Secchi's original classes I-II-III. Today's types L, T, and Y are mostly brown dwarfs, substellar objects that seem to form, like stars, by gravitational collapse, but are not massive enough to get hot enough to do sustained hydrogen fusion. Secchi's carbon stars are mostly giants and have their own nomenclature; some stars, including some especially hot stars, show emission lines in their spectra and would correspond to Secchi's class V.

Additional codes specify distinctive or peculiar features of an individual spectrum. By way of example: an Am star would indicate an A star with enhanced metal lines; a Be star would indicate a B star with prominent emission lines. The following sketch provides a graphical representation of approximately how the strengths of lines of several prominent ionization species vary across spectral types. The subsequent table lists a few prominent features used to define the various spectral types.

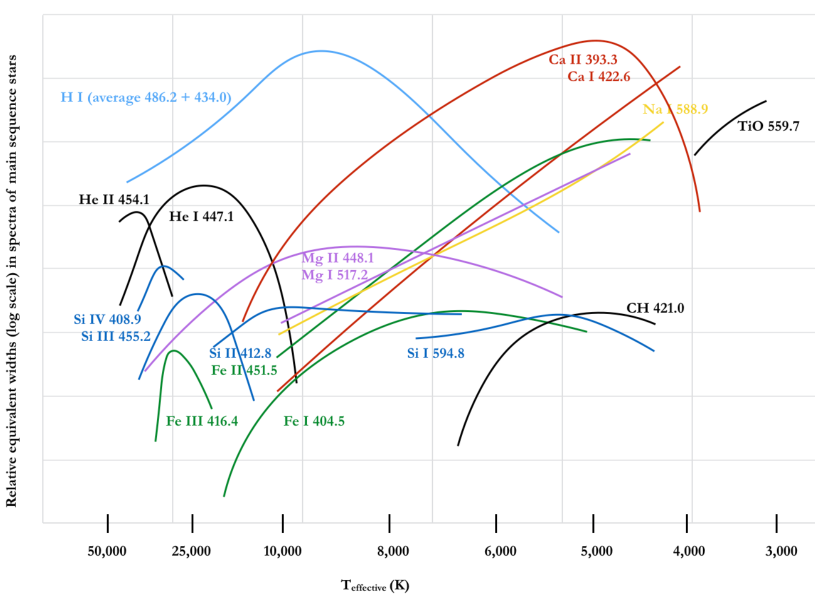

Figure 14.3: Line strengths as a function of effective temperature for main sequence stars of roughly solar composition.

Based on data collected in C. Jaschek and M. Jaschek, *The Behavior of Chemical Elements in Stars*

Table 14.1: some of the principal characteristics of the spectral types.

O	blue	T_{eff} >30,000 K	He II; multiply ionized metals; some lines in emission
B	blue-white	T_{eff}: 10,000 - 30,000 K	He I, Balmers increasing
A	blue-white - white	T_{eff}: 7,500 - 10,000 K	Balmer lines; ionized metals; Ca II strengthening
F	white - yellow-white	T_{eff}: 6,000 - 7,500 K	Balmer lines ↓; Ca II ↑; neutral metals increasing
G	yellow-white - yellow	T_{eff}: 5,200 - 6,000 K	Balmers weak; Ca II max; neutral metals ↑; CH
K	yellow - orange	T_{eff}: 3,700 - 5,200 K	Balmers very weak; Ca II; neutral metals strong; TiO
M	orange - red	T_{eff}: 2,400 - 3,700 K	Ca II ↓; neutral metals strong; molecular bands

And a few more spectral types:

WR	Wolf-Rayet stars; no hydrogen lines; emission lines of helium and N &/or C, sometimes O
L	mostly brown dwarfs; some very low-mass stars; some very cool supergiants
T	infrared brown dwarfs, T_{eff} ~ 700 - 1,300 K; methane present in spectra
Y	recently created classification for very cool brown dwarfs; possibly NH_3, H_2O present in spectra
C	Carbon stars; most are CR, CN giants similar to G/K and K/M stars but with added carbon
S	ZrO in spectra; carbon intermediate between carbon stars and normal M stars
wd or D	white dwarfs: degenerate remnants of solar-type stars; broad lines

An excellent resource for representative stellar spectra is a 1992 article by David R. Silva and Mark E. Cornell: “A New Library of Stellar Optical Spectra”, *Ap.J. Supplement Series*, **81**, pp. 865 - 881. The following plot shows several of these spectra from 350 to 680 nm; they are offset vertically for clarity by an arbitrary amount in flux. Notice how the peak shifts to longer wavelengths for the cooler spectral types and that the Balmer lines are strongest in early A stars.

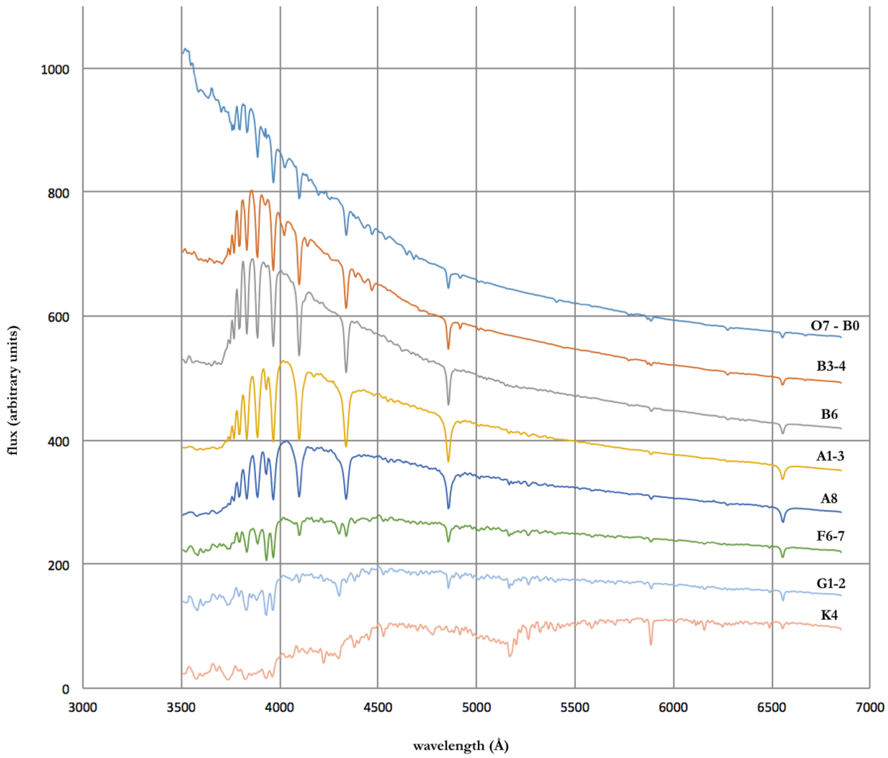

Figure 14.4: Representative spectra for main sequence stars.

Based on data from Silva and Cornell 1992, *ApJSS* **81,** 865.

One of the first to notice the distinction between the line widths of spectra of stars such as the Sun and giants was Antonia Maury, another of the women of the Harvard College Observatory. Maury was a bit ahead of her time, proposing a somewhat complicated spectral classification scheme in the decades before an adequate understanding of electron energy levels existed. When astronomers did adopt a luminosity classification system to distinguish giant and supergiant stars, for some perverse reason we decided to use Roman numerals. . .which have nothing whatsoever to do with Secchi's original use of Roman numerals. The Sun is a G2 V star, meaning spectral type G2 and luminosity class V, placing it among the hydrogen-fusing stars. Red giant stars are luminosity class III, supergiants class I. To make more sense of this let's turn to an important graphical way of representing the parameter space of stellar colors and magnitudes.

The Hertzsprung - Russell and color-magnitude diagrams

In the early 1900s Ejnar Hertzsprung and, independently, Henry Norris Russell were two of the first astronomers to plot stars' spectral types vs. their absolute magnitudes and to realize that stars are not randomly distributed on such a plot. Today such a plot is often simply called an *H-R diagram*. The H-R diagram plotted below shows some of the brightest and the nearest stars in our sky.

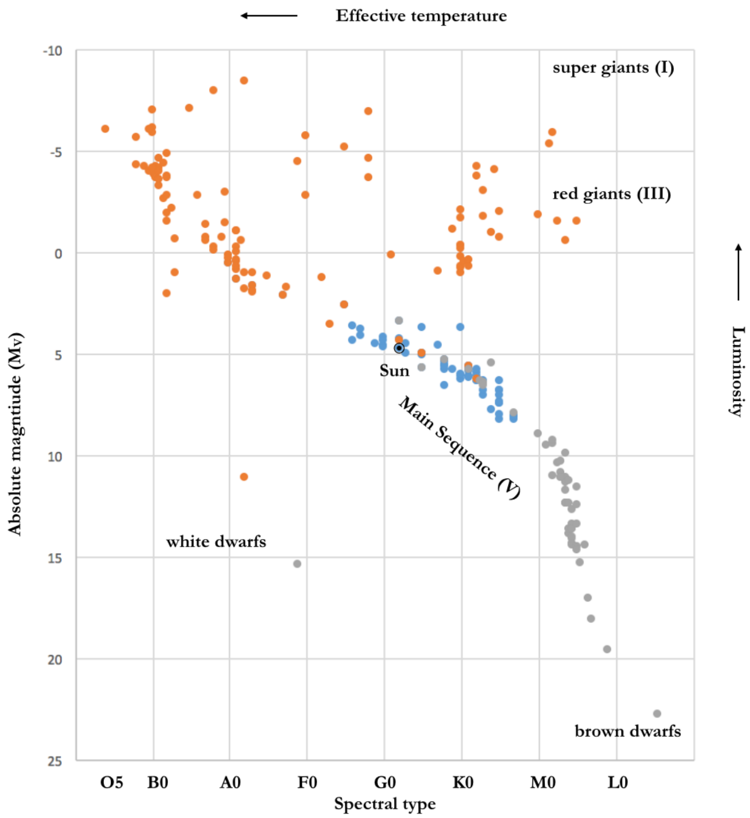

Figure 14.5: H-R diagram.

Orange: brightest stars in our sky;
Blue: nearby (≲10 pc) bright stars;
Gray: nearest (≲8 pc) stars;
plus 2 wd companions.

The Sun is a G2 star, M_V ~4.8.

There are several important points to notice here. First, the Main Sequence: these are the stars doing hydrogen fusion (tailing off at the faint end into the brown dwarfs). Second, the majority of nearby stars are fainter than the Sun. Most of these stars would not be visible very far away; if this is indicative of the general population of stars elsewhere in our galaxy or in other galaxies, it suggests that the majority of all stars are smaller and fainter than the Sun. The brightest appearing stars in our sky are thus not representative of stars in general. The brightest appearing stars in our sky are more likely to be rare, very luminous stars, visible for great distances. The Roman numerals — V, III, I — indicate the major *luminosity classes*, mentioned above. Supergiants are subdivided into luminosity class Ia and Ib and there's a class II, called bright giants. Subgiants, stars falling between the main sequence and the red giant branch, are class IV. The hydrogen-fusing main sequence stars, class V, are sometimes collectively called dwarf stars, to distinguish them from the various types of giant stars. Some stars, especially those with lower metallicity, fall just below the Main Sequence and are referred to as subdwarfs, luminosity class VI. Stellar remnants such as white dwarfs are not assigned a luminosity class in this scheme. We will consider in the stellar evolution chapter how stars come to find themselves in the various regions of the diagram.

We mentioned the pressure broadening of spectral lines, above. Stars of the same spectral type have the same strength ratios; e.g., the strength of the Ca II K line relative to Hγ would be the same. Stars of a given spectral type but different luminosity class would have the line strength ratios appropriate for their spectral type but overall the lines would be broader for main sequence stars and skinnier for giants. Stellar radii are largest for stars in the upper right-hand corner of the H-R diagram — recall the Stefan-Boltzmann equation: $L = 4\pi R^2 \sigma T^4$. Effective temperature increases toward the left, luminosity increases upwards, and that means radius must increase diagonally from lower left to upper right. If a star is going to be cool and luminous, it must be huge. Gravitational acceleration $g = {GM}/{r^2}$ and is going to be lower for a very distended star with a huge radius. Thus the pressure broadening effects can also be called surface gravity effects. Stars with lower luminosity class, e.g., the type I supergiants, have lower surface gravities as well.

The stars plotted in the H-R diagram above are relatively similar to the Sun in the compositions of their photospheres. Fusion reactions are changing the core compositions but, for almost all stars, the products of those fusion reactions are not dredged up to the surface and the composition that we would measure from the light received from the photosphere doesn't directly tell us about the composition changes in the star's core. Photospheric compositions may change slightly over the lifetime of a star as heavy elements may slowly settle out. Stars such as the Sun or those that have formed more recently have formed from material that has been noticeably enriched in heavy elements. Astronomers use *X, Y, Z* to represent the mass fraction of H, He, and heavy elements, the latter collectively called "metals". Stars with compositions roughly similar to the Sun have *Z* values on the order of 1-2%. Stars such as these are called Population I. Stars that are much older, having formed perhaps 12-13 billion years ago, formed at a time when the universe was much less enriched in heavy elements. These stars may have *Z* values 10 - 100 times lower than the Sun. Such low-*Z* stars are called Population II. For completeness, the earliest stars in the universe, which would have formed from almost nothing beyond hydrogen and helium, are called Population III.

Terminology note: we often use the term *metallicity* to refer to *Z*, this heavy-element mass fraction. More specifically, astronomers may use metallicity to refer to a measure of a star's iron content. Iron is relatively abundant and has many absorption lines in the visible part of the star's spectrum, which together make its abundance relatively easier to determine. Iron content thus becomes a proxy for metal abundance in general. Metallicity in this context refers to the logarithm of a star's iron-to-hydrogen ratio as a fraction of the Sun's iron-to-hydrogen ratio:

$$[\mathrm{Fe/H}] \equiv \log_{10}\left(\frac{\left(N_{\mathrm{Fe}}/N_{\mathrm{H}}\right)_{star}}{\left(N_{\mathrm{Fe}}/N_{\mathrm{H}}\right)_{Sun}}\right).$$

N in this context refers to number density. An [Fe/H] value of −2 means an iron abundance that is only 1/100 the solar iron abundance. If we want to be a bit more precise, using iron abundance as a proxy is a good first approximation, but it's only an approximation. The evolution of the abundances of the various heavy elements likely did not proceed in lockstep because different groups of elements are produced by different processes involving stars of different masses; the evolution of those stars don't all proceed at the same rate, and thus neither can the early enrichment of all elements heavier than H and He. It's thought that elements such as C, N, O get a head start in enrichment, being the most likely to have been produced by the very first, massive, short-lived stars. This is supported by evidence for the existence of very old, very low mass stars which show some C, N, O but still no iron. (These are called CEMP stars, which stands for carbon-enhanced metal-poor.)

The horizontal axis in the H-R diagram is principally an indicator of surface temperature. We could plot effective temperature rather than spectral type or we could plot color, often $(B-V)$. Both spectral type and color have their observational advantages: color indices may take less time to acquire at the telescope but suffer from reddening effects for stars at large distances; spectral types may take more time to acquire, because spreading a star's light out into a spectrum means that less light falls on any given portion of our detector surface, but suffer much less from dust. Plotting color is very useful for stars of a known distance, or at least for stars known to be all at basically the same distance, as in members of a star cluster.

Absolute magnitude or luminosity plotted on the vertical axis allows us to compare stars of different distances, but, as with color, if we are interested in the distribution of stars from a single cluster, there's no reason not to plot apparent magnitude. In the *color-magnitude diagram*, or CMD, we often plot $(B-V)$ color on the x-axis and *V* magnitude on the y-axis, although we are of course not limited to using the *B* and *V* filters.

As we'll see when we discuss stellar evolution, cluster color-magnitudes are very useful for understanding the age and composition of star clusters. Stars that are forming, or have formed relatively recently, in our Milky Way galaxy form in *open* (or *galactic*) *clusters*. These tend to be groups of several hundred stars each, and include clusters such as the Pleiades or the Hyades (the head of Taurus). There are a few open clusters that have ~ 10^4 stars. Open clusters tend to lie in or very near the plane of the galaxy and over time, as they move in and out of spiral arms and experience various gravitational nudges from passing stars, the stars in open clusters will drift apart. For example, in 2021 researchers analyzing precise stellar position and motion data from the Gaia spacecraft reported identifying a tail of stars that have been pulled out from the Hyades cluster.

The oldest stars clusters in our galaxy are *globular clusters*, objects such as M13 in Hercules and M15 in Pegasus. Globular clusters often contain over 10^5 stars, these being older, low-metallicity, Population II stars. The color-magnitude diagrams for open and globular clusters tend to look rather different. One reason is the age difference. High-mass stars live and die faster than their low-mass siblings, meaning that as clusters age they will tend to lose high-mass stars from the main sequence as those stars change to become giants or supergiants. Another reason is the composition difference. Metals often have absorption lines in the blue and green and, more generally, metals contribute to the overall opacity in a star (e.g., a photon might go into ionizing a metal atom rather than simply escaping from a star). This means that low-Z stars are going to be bluer and a tad more luminous than high-Z stars for a given mass.

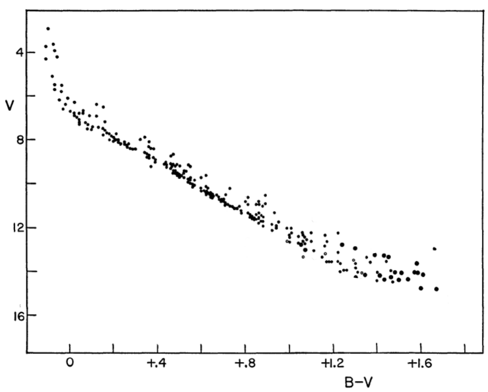

Figure 14.6a: Pleiades CMD. From Johnson & Mitchell 1958, *ApJ* **128**, 31.

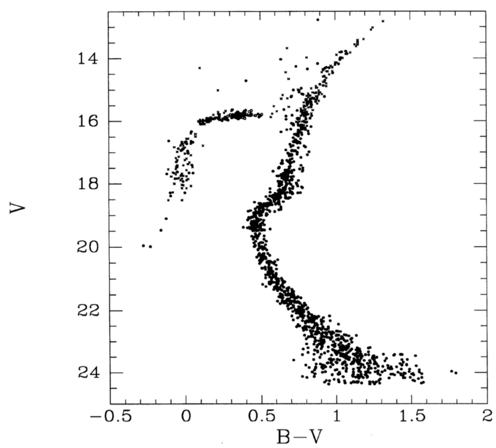

Figure 14.6b: M15 CMD. From Durrell & Harris 1993, *AJ* **105**, no. 4, 1420.

The two figures above show representative color-magnitude diagrams for these two types of clusters, in this case the Pleiades, an open cluster roughly 1% solar age, and the globular cluster M15. The Pleiades still has relatively massive stars on the Main Sequence; in M15 stars at the top of the Main Sequence have evolved off and there is a well developed distribution of red giants. (Star clusters are also discussed in the chapter on the Milky Way; images of these two are below.)

Acknowledging the fact that ages and compositions and dust along the line of sight are different, it's still very clear that M15 is more distant than the Pleiades. If Main Sequence stars of a given $(B - V)$ are roughly similar in absolute magnitude then we can use the offset between the Main Sequences of these two clusters to estimate their relative distances. For instance, a star with $(B - V) = 0.6$ has an apparent magnitude m_V ~10 in the Pleiades and ~21 in M15. From the distance modulus equation we have that $m_1 - m_2 = 5\log(d_1 / d_2)$ for two stars of the same M_V and extinction. For a magnitude difference of ~11, that's a distance ratio of ~150, which is a bit of an overestimate (~100 would be better).

If we properly account for extinction and compare Main Sequence stars of similar metallicities we can use the offset $(m - M)$ between observational and theoretical Main Sequences to estimate a cluster's distance. Depending on whether the observations are of spectral types or colors, this technique is called spectroscopic or photometric parallax. That's a bit of a misnomer, because it doesn't involve observations of parallax, although since parallax and distance are simply inverses of each other, it's not all that misleading.

In the next chapter we will talk about the lives of stars and pay particular attention to what we can learn about stellar evolution from studying H-R and CM diagrams.

Figure 14-7a

Left: Pleiades open cluster. Credit: NASA, ESA and AURA / Caltech Palomar Observatory. http://hubblesite.org/image/1562/news_release/2004-20

Right: M15 globular cluster. Credit: NASA, ESA. http://www.spacetelescope.org/images/heic1321a/

Figure 14-7b

Stellar properties

Table 14.2: approximate properties of Population I main sequence (luminosity class V) stars

	T_{eff}	B−V	M_V	B.C.*	$R / R_\odot$	$M / M_\odot$
O2	48,000	−1.0	−6.1	−4.7	15	100
O5	42,000	−0.33	−5.7	−4.40	12	60
B0	30,000	−0.3	−4.0	−3.16	7.4	17.5
B5	15,200	−0.17	−1.2	−1.46	3.9	5.9
B8	11,400	−0.11	−0.25	−0.80	3.0	3.8
A0	9,790	−0.02	0.65	−0.30	2.4	2.9
A5	8,180	0.15	2.0	−0.15	1.7	2.0
F0	7,300	0.30	2.7	−0.09	1.5	1.6
F5	6,650	0.44	3.5	−0.14	1.3	1.4
G0	5,940	0.58	4.4	−0.18	1.1	1.05
G5	5,560	0.68	5.1	−0.21	0.92	0.92
K0	5,150	0.81	5.9	−0.31	0.85	0.79
K5	4,410	1.15	7.4	−0.72	0.72	0.67
M0	3,840	1.40	8.8	−1.38	0.60	0.51
M5	3,000	1.64	12.3	−2.73	0.2	0.15
M8	2,400	2.12	18.7	−4.1	0.10	0.08

*B.C. = bolometric correction

Table 14.3: roughly, approximate properties of Population I supergiants (luminosity class I)

	T_{eff}	B−V	M_V	B.C.	$R / R_\odot$
O9	32,000	−0.27	−6.5	−3.2	25
B2	17,600	−0.17	−6.4	−1.6	40
B5	13,600	−0.10	−6.2	−0.95	50
A0	9,980	−0.01	−6.3	−0.4	60
A5	8,610	0.09	−6.6	−0.1	65
F0	7,460	0.17	−6.6	−0.01	80
F5	6,370	0.32	−6.6	−0.03	100
G0	5,370	0.76	−6.4	−0.15	120
G5	4,930	1.02	−6.2	−0.3	150
K0	4,550	1.25	−6.0	−0.5	200
K5	3,990	1.60	−5.8	−1.0	400
M0	3,620	1.67	−5.6	−1.3	500
M5	2,880	1.80	−5.6	−3.5	800

Table 14.4: properties of a few interesting and/or well-known stars.

	RA (2000)			Dec (2000)		m_B	m_V	B-V	parallax	μ RA	μ Dec	v_{radial}	Spt& LC	notes
	h	m	s	°	′				mas*	mas	mas	km/s		
α Cen A	14	39	36.5	-60	50.0	0.72	0.01	0.71	754.81	-3679	474	-21.4	G2V	
α Cen B	14	39	35.1	-60	50.3	2.21	1.33	0.88	796.92	-3614	803	-20.7	K1V	
α Cen C	14	29	42.9	-62	40.8	12.95	11.13	1.82	769.8	-3776	766	-22.4	M5.5Ve	+ planet
Albireo (β Cyg) A	19	30	43.3	27	57.6	4.17	3.09	1.08	7.51	-7.2	-6.2	-24.07	K3II+B9.5 V	
Albireo (βCyg) B	19	30	45.4	27	57.9	5.01	5.11	-0.1	8.38	-0.99	-0.5	-18.80	B8Ve	
Aldebaran α Tau)	4	35	55.2	16	30.6	2.4	0.86	1.54	48.94	63.5	-189	54.26	K5III	
Algol (β Per)	3	8	10.1	40	57.3	2.07	2.12	-0.05	36.27	2.99	-1.7	4.0	B8V	eclipsing binary
Altair (αAql)	19	50	47	8	52.1	0.98	0.76	0.22	194.95	536	385	-26.6	A7Vn	
Antares (α Sco)	16	9	24.5	-26	25.9	2.75	0.91	1.84	5.89	-12	-23	-3.5	M0.5Iab+ B3V	
Arcturus (α Boo)	14	15	39.7	19	10.9	1.18	-0.05	1.23	88.83	-1093	-2000	-5.19	K1.5III	
Barnard's	17	57	48.5	4	41.6	11.24	9.51	1.73	547.5	-803	10362.54	-110.51	M4V	BY Dra var
Betelgeuse (α Ori)	5	55	10.3	7	24.4	2.27	0.42	1.85	6.55	27.5	11	21.91	M1-2Ia-Iab	
Canopus (α Car)	6	23	57.1	-52	41.7	-0.59	-0.74	0.15	10.55	19.9	23.2	20.3	A9II	
Capella (α Aur)	5	16	41.4	45	59.9	0.88	0.08	0.8	76.2	75.3	-427	29.19	K0+G1III	spectroscopic bin
Castor (α Gem) Aab	7	34	35.9	31	53.3	1.62	1.93	-0.31	64.12	-206.3	-148	6	A1.5IV+d M1e	spectroscopic bin
Castor (α Gem) Bab	7	34	36.1	31	53.3	1.62	2.97	-1.35	64.12	-206.3	-148	-1.2	A1IV+dM 1e	spectroscopic bin

Castor (α Gem) Cab	7	34	37.6	31	52.2	10.56	9.27	1.29	64.12	-207.6	-96	2.5	2(M0.5Ve)	BY Dra &bin
τ Ceti	1	44	4.1	-15	56.2	4.22	3.5	0.72	273.96	-1721	854	-16.68	G8V	multiple planets; debris disk
61 Cygni A	21	6	53.9	38	45.0	6.39	5.21	1.18	285.95	416	32	-65.74	K5V	BY Dra var
61 Cygni B	21	6	55.3	38	44.5	7.4	6.03	1.37	286.15	4106	3156	-64.07	K7V	
Deneb (α Cygni)	20	41	25.9	45	16.8	1.34	1.25	0.09	2.31	2.0	1.9	-4.9	A2Ia	
ε Eri	3	32	55.8	-9	27.5	4.61	3.73	0.88	310.9	-975.2	19.5	16.43	K2 V	BY Dra var; + planet
Fomalhaut (α PsA) A	22	57	39	-29	37.3	1.25	1.16	0.09	129.8	329	-165	6.5	A4V	+planet; debris disk; angular size 0.212 ′
Fomalhaut B (TW PsA)	22	56	24.1	-31	33.9	7.58	6.48	1.1	131.4	330	-158	7.2	K4Ve	BY Dra var
Fomalhaut C	22	48	4.5	-24	22.1	14.3	12.624	1.676	130.3	332	-184	6.5	M4.0Ve	
ε Indi	22	3	21.7	-56	47.2	5.75	4.69	1.06	274.8	3967	-2536	-40	K5V	
Luten 726-8 A (BL Cet)	1	39	1.5	-17	57.0		12.7		374	3296	563	29	M5.5V	flare
Luten 726-8 B (UV Cet)	1	39	1.5	-17	57.0		13.2		374	3296	563	29	M6V	flare
Mizar (ζ 1 UMa) A	13	23	55.5	54	55.5	2.29	2.23	0.06	38	121	-22	-5.6	A1.5V	spectroscopic bin
Mizar (ζ 2 UMa) B	13	23	56.3	54	55.3	4.05	3.88	0.17	40.5	114	-26.5	-9.3	A1+A7IV-V	spectroscopic bin
Alcor (80 UMa)	13	25	13.5	54	59.3	4.18	4.01	0.17	40.5	113	-28.6	-8.9	A5+M3-4 V	binary
Polaris (α UMi) Aa	2	31	49.1	89	15.9	2.62	2.02	0.6	7.54	44.5	-11.9	-16.4	F8 Ib	Cepheid
Polaris (α UMi) Ab							9.2						F6V	
Polaris (α UMi) B	2	30	33.5	89	15.6	8.69	8.2	0.49				-8	F3V	
Pollux (β Gem)	7	45	18.95	28	1.6	2.14	1.14	1.0	94.54	-627	-45.8	3.23	K0IIIb	+ planet
Procyon (α CMi) A	7	39	18.1	5	13.5	0.79	0.37	0.4	284.6	-717	-1035	-3.2	F5 IV-V	
Procyon (αCMi) B	7	39	17.9	5	3.4	10.7	10.92	-0.2		-709	-1024		DQZ	
R136a1	5	38	42.4	-69	6.0	12.78	12.77	0.01	0.02				WN5h	LMC
Regulus (α Leo) A	10	8	22.3	11	58.0	1.24	1.4	-0.2	41.13	-249	5.6	5.9	B8IVn	
Regulus (α Leo) B/C	10	8	12.8	11	59.8	8.99	8.13	0.9	41.21	-254	8.0	6.72	K0Ve (+M4V?)	binary
Rigel (βOri) A	5	14	32.3	-8	12.1	0.1	0.13	-0.03	3.78	1.3	0.5	17.8	B8Iae	
Rigel (βOri) Bab	5	14	32	-8	12.2	10.4	10.4	0.00				19.1	B9+B9	spectroscopic bin
Ross 128	11	47	44.4	0	48.3	12.905	11.153	1.75	295.8	607	-1223	-31.2	M4V	flare; +planet
Sirius (α CMa) A	6	45	8.9	-16	43.0	-1.09	-1.09	0.00	379.2	-546	-1223	-5.5	A1V	spectroscopic bin
Sirius (α CMa) B	6	45	9	-16	43.1	8.41	8.44	-0.03		-547	-1207		DA1.9	
Spica (α Vir)	13	25	11.6	-11	9.7	0.74	0.97	-0.23	13.06	-42	-30.7	1	B1V	
Vega (αLyr)	18	36	56.3	38	47.0	0.03	0.03	0.00	130.23	201	286	-20.6	A0V	
WISE 0855-0714	8	55	10.8	-7	14.7				448	-4800	500		Y2	m_J = 25

* mas means milli-arc-seconds.

Most data in this table are taken from the SIMBAD database / Strasbourg astronomical Data Center; the latter is located at: http://cds.u-strasbg.fr

Sample problems

1. The distance to Deneb is a bit uncertain, in part because it has such a bright apparent magnitude; a possible value for its parallax, at least as of 2018, is 2.31 mas.

a) If the radius of Deneb is ~203 solar radii and its effective temperature is 8525, calculate Deneb's luminosity in units of solar luminosities.

b) If the bolometric magnitude of Deneb is ~ −8.49, calculate Deneb's luminosity in units of $L_{\odot}$. The bolometric magnitude of the Sun is 4.75.

2. Suppose that the parallax for a main sequence star has a measurement uncertainty such that the error in distance could be 7%. What effect would that have on the location of the star on the H-R diagram? You can ignore changes in extinction.

3. Explain how and why the strengths of the Balmer lines vary across the spectral types.

4. Describe several line-broadening mechanisms.

5. Why aren't absorption lines in stellar spectra totally black in the center; i.e., why isn't the line center flux = 0?

6. Reading carefully? Explain / define

a) effective temperature
b) opacity
c) equivalent width
d) ionization energy
e) metallicity
f) brown dwarf

Answers to selected problems are on the next page:

1. a) This is a $4\pi r^2\sigma T^4$ ratio question; $L_D \sim 198{,}000\ L_\odot$

   b) This is a $L_1/L_2 = 10^{0.4(M_2-M_1)}$ ratio question; $L_D \sim 195{,}000\ L_\odot$

2. We'd be off in absolute magnitude by ~0.15.

3. Think about the effective temperatures of stars of various spectral types and the peak wavelengths of their spectra, and hence energy available to excite or ionize hydrogen. Recall that the Balmer absorption lines arise from the n=2 energy level, meaning that for a star's spectrum to have strong, i.e., very dark, Balmer absorption lines there must be a high probability that the hydrogen atoms' electrons will be in the n=2 level and not either stuck down in the ground state or removed from the proton all together.

5. The short answer is that the surface temperature of the star is not 0.