1

Background

Starlight tells us about the nature of stars. Spectroscopy is the key by which the starlight is decoded. From stellar spectra, both lines and continua, we extract the fundamental stellar properties of surface gravity, temperature, and chemical composition. From there we can often deduce the mass, the most fundamental property of all. But there is much more. In the spectral lines, we see signatures of the velocity fields pervading stellar atmospheres, which in turn uncover details of surface convection and the mottled appearance of stellar surfaces. Chemical markers tell us about deep convection. The broadening of spectral lines tells us the rotation rates of stars, and sometimes time variations of lines allow us to map spots on the surface. Stars couple convection and rotation to generate magnetic fields. Magnetic fields, together with winds of escaping material, dissipate angular momentum, and this can be studied using the measured rotation rates. Knowledge of chemical composition informs us about nuclear processes, energy generation, and element building. Differences in chemical composition help us map out the dynamical structure of the Milky Way and teach us about star formation processes. And the list goes on. Stellar spectroscopy is the connecting link between the observations and the rest of stellar astrophysics.

The topics brought together in this chapter are background material to set the stage.

What Is a Stellar Photosphere?

The photosphere is the region of a star's atmosphere from which the major portion of the light comes. The whole atmosphere is a transition region from the stellar interior to the interstellar medium. One way to quantify this description is to look at the change in average kinetic temperature with height, the Sun being a close-at-hand example. Figure 1.1 shows the solar temperature profile with the four basic sections labeled, sub-photosphere, photosphere, chromosphere, and corona. The visible spectrum comes from the photosphere. When we speak of the size of a star, for example its radius, we mean the size of the apparent stellar disk as defined by the photosphere. When we talk about a star's temperature, we fundamentally mean the characteristic temperature of the photosphere. And if we think about the surface of a star, it is the photosphere we have in mind. Light from the sub-photospheric layers does not penetrate to the surface, so we cannot see those layers. The outer layers, the chromosphere and the corona, emit most of their light in the ultraviolet (UV) and x-ray regions of the spectrum.

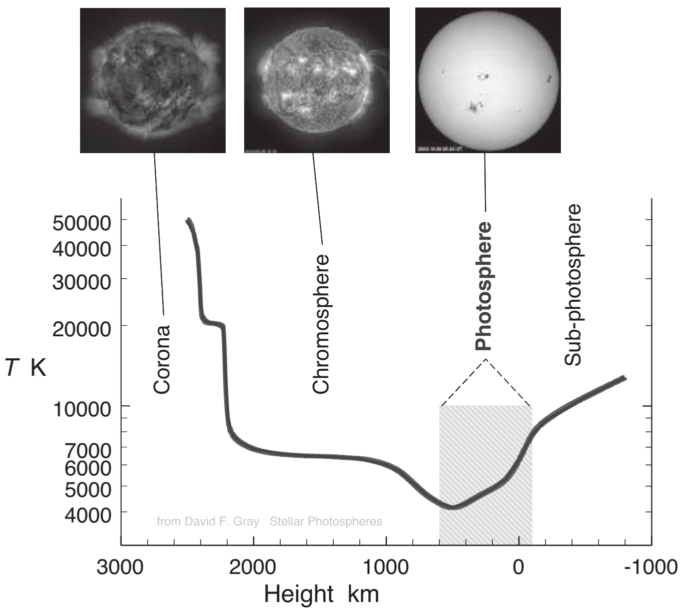

Figure 1.1 The temperature in the outer layers of the Sun is shown as a function of geometrical height in kilometers. The center of the Sun is off to the right. Visible light comes from the photosphere. Based on Vernazza *et al*. (1973). Images, left: inner corona Fe IX/X 171 Å line; center: chromosphere at 304 Å; right: photosphere in white light with large sunspots. Images courtesy of SOHO consortium (ESA & NASA).

The solar photosphere is ~700 km thick, as illustrated in Figure 1.1. It is only by virtue of its small extent compared to the much larger solar radius, $R_\odot = 696\,100$ km, that the edge of the Sun appears sharp. The circumpunct, ⊙, is used to denote solar values. The geometrical extent of the photosphere in other stars differs inversely with the *surface gravity* defined by

$$g = g_\odot \frac{\mathfrak{M}}{R^2}, \tag{1.1}$$

where $g_\odot = 2.740 \times 10^4$ cm/s^2; $\log g_\odot = 4.438$ is the surface gravity of the Sun, and the mass, $\mathfrak{M}$, and radius, R, of the star are in solar units (see Appendix A). In astronomical culture, centimeter–gram–second (cgs) units are frequently used as the default unit system. This is especially true for surface gravity numbers, where $\log g$ is assumed to be in cm/s^2 units.

The overall thickness of the photosphere depends upon the opacity of the gases forming it. In cool stars, the opacity is larger and the photosphere thinner when the gas is richer in elements other than hydrogen and helium. In the somewhat arcane astronomical terminology, these elements are called “metals.”

The third physical variable strongly affecting the nature of the atmosphere is its characteristic temperature. Typically the temperature drops by somewhat more than a factor of two from the bottom to the top of the photosphere, as in Figure 1.1. Instead of choosing a temperature at some depth to characterize the temperature of a star, it is customary to use the *effective temperature*, T_{eff}. The effective temperature is defined in terms of the total power per unit area radiated by the star,

$$\int_0^\infty \mathfrak{F}_\nu \, d\nu \equiv \sigma T_{\text{eff}}^4, \tag{1.2}$$

where the total radiant power per unit area is given by the integral, and $\sigma = 5.6704 \times 10^{-5}$ erg/(s cm^2 K^4) is the Stefan–Boltzmann constant. Here $\mathfrak{F}_\nu$, or alternatively $\mathfrak{F}_\lambda$, is the flux leaving the stellar surface and is formally defined in Chapter 5. For now, think of it as "the amount of light" in units of erg/(s cm^2 Hz) or erg/(s cm^2 Å). Equation 1.2 has the form of the Stefan–Boltzmann law, which we will meet in Chapters 5 and 6, making T_{eff} the temperature of a black body having the same power output per unit area as the star. However, the distribution of power across the spectrum may differ dramatically from a black body at the same effective temperature.

The total power emitted by the star is termed the *luminosity*. Luminosity is related to the effective temperature through the surface area, namely,

$$L = 4\pi R^2 \int_0^\infty \mathfrak{F}_\nu \, d\nu = 4\pi R^2 \sigma T_{\text{eff}}^4 \tag{1.3}$$

for a spherical star of radius R.

Traditionally and historically, the more empirical parameters of spectral type, magnitude, and color index have been used to describe stellar temperature and surface gravity.

Spectral Types

The most powerful way to analyze starlight is through spectroscopy. The first step in many cases is to classify a star by spectral type. Thanks to many studies done over past decades, knowing the spectral type tells us what kind of star we are dealing with, hot or cool, big or small, and so on. The concept of a spectral type is to classify stars into groups according to the strengths of their spectral lines. The lines in stellar spectra are overwhelmingly seen as absorption lines, i.e., regions where light is missing. The strongest of these were first seen in the solar spectrum by Joseph Fraunhofer (1787–1826) and are sometimes referred to as "Fraunhofer lines." Some of these are listed in Appendix D with their Fraunhofer letter designations, many of which are still in common use, such as the sodium D lines.

The format of a spectral type consists of a capital letter followed by an Arabic number representing the fraction toward the next letter class, followed by a Roman numeral. The letter is the temperature classification, the Roman numeral the luminosity classification. Spectrograph dispersions of ~100 Å/mm, giving ~2 Å resolution, are best for classification. Higher resolution resolves details of line profiles giving unwanted complications; lower resolution does not adequately separate individual lines. The traditional approach uses photographic records, called spectrograms, that are centered in the blue region of the

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spectrum where photographic emulsions are most sensitive, but modern electrical detectors work equally well or better. Spectrograms are in black and white, rather than in color, because color photography is much less sensitive to the faint starlight. But the position or wavelength of a line essentially tells us the color, so color photography would be redundant.

In the traditional approach, line strengths and ratios are estimated visually by comparing an untyped spectrogram with standards. The more modern approach uses graphs of the spectra. Establishing a grid of standard stars is a major undertaking with deep historical roots (see Huggins & Huggins 1899, Draper 1935, Garrison 1984, Hearnshaw 1986, Keenan 1985, 1987). Figure 1.2 shows a set of standard spectra in order of decreasing effective temperature. The height of the spectrum has no meaning; it just helps us see the spectral lines. Those spectra toward the cool end of the sequence are referred to as "late-type" spectra, and those on the hot end as "early-type" spectra. In principle the spectral classification separates stars into about 60 temperature steps. These range (hottest to coolest) from O2 to M8 through the standard (temperature) sequence: O, B, A, F, G, K, M with 10 subdivisions per letter group. But some subclasses are rarely used. For example, the main subtypes for the cooler stars are G0, G2, G5, G8, K0, K1, K2, K3, K4, K5, M0, M1, M2, M3, and M5. In a few cases, half intervals are given full weight as a subclass, for example, O9.5 and B2.5. Statistically smooth relations exist between these main subtypes

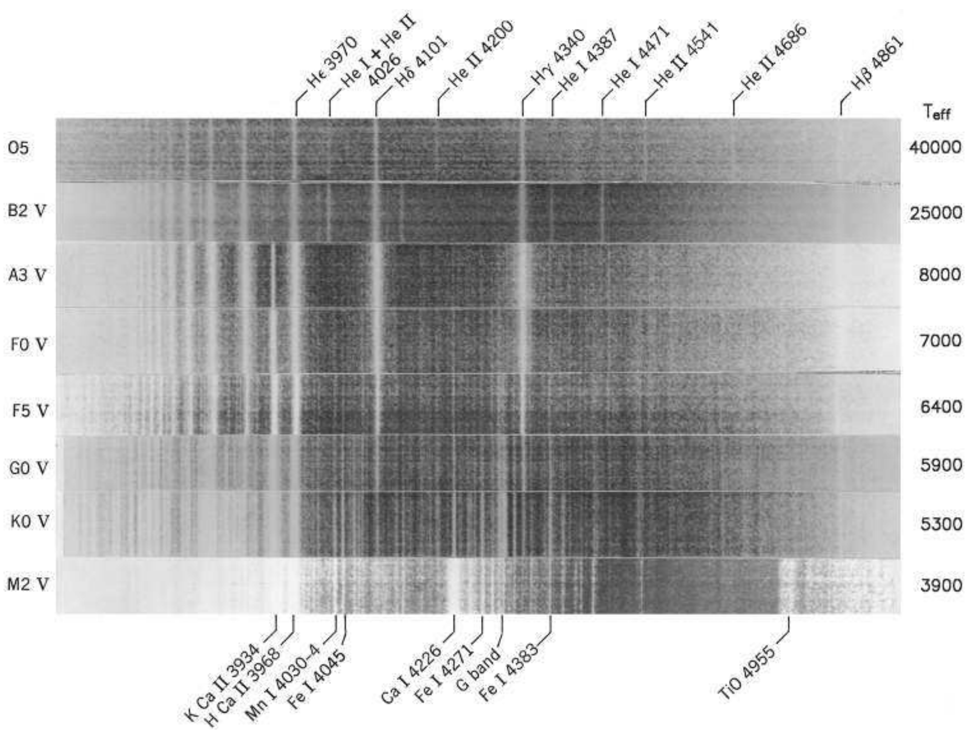

Figure 1.2 The temperature classification sequence is illustrated by these standard spectral types from Abt *et al*. (1968). Notice how the hydrogen lines reach maximum absorption in the A type and how most lines increase in strength toward cooler stars. Specific lines used for classification are labeled. Direct photographs of spectra are "negatives," meaning light is dark and vice versa.

and other temperature indicators such as color indices (to be discussed shortly). The Sun's spectral type is G2 V by definition. The effective temperature associated with each spectral type requires a separate calibration, and is discussed in Chapter 14.

Spectral classification of very cool stars includes extension to classes L (2200–1300 K) and T (1300–800 K), and on down to planet temperatures (see Geballe *et al*. 2002, Leggett *et al*. 2002). There are also side branches: R stars parallel K stars, but show carbon molecular bands; N stars parallel M stars, but show very strong carbon molecular bands; S stars parallel M stars, but have the usual titanium oxide bands overshadowed by those of zirconium oxide (see Hearnshaw 1986, Gray & Corbally 2009).

Luminosity-sensitive features in the spectrum show markedly less variation, and only a few classes are recognized: 0, I, Ia, Ib, II, III, IV, and V in order of decreasing luminosity. The names associated with these numerals are 0 hypergiants, I supergiants, II bright giants, III giants, IV subgiants, and V dwarfs. Sometimes subdwarfs are denoted as class VI. Luminosity effects are illustrated in Figure 1.3 and their physical causes are discussed in later chapters. Sometimes the suffixes listed in Table 1.1 are used to supply additional information.

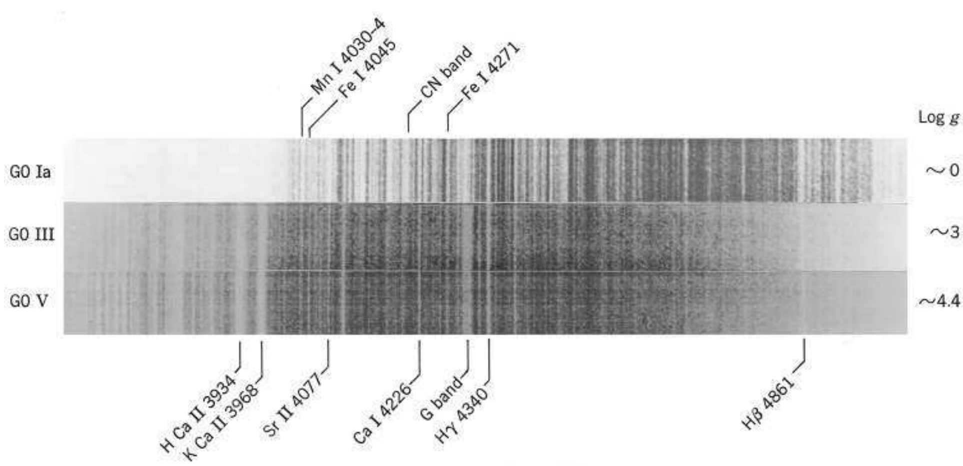

Figure 1.3 Luminosity classification is illustrated by these standard spectra from Abt *et al*. (1968). Specific lines used for classification are labeled.

A summary of criteria, based on Morgan *et al*. (1943), Keenan and McNeil (1976), and Morgan *et al*. (1978) is given in Table 1.1. Before classification can begin, the observer must acquire spectrograms of the standard stars using the same equipment and technique used with the program stars. The first step is to identify the lines whose strengths serve as the criteria for classification. We easily recognize the hydrogen lines from class O to F5, and the less conspicuous lines can be referred to them. As seen in Figures 1.2 and 1.3, H_β remains distinguishable as the most intense line at the long-wavelength end of the spectrogram as late as K0. From F5 to M, the G band near the center of the spectrogram and the H and K lines on the short-wavelength end can be used as guideposts.

The spectrum is first classed approximately, following the flow chart in Figure 1.4. Does it have strong hydrogen lines, many lines, or does it show molecular bands? This rough

Table 1.1 *Spectral classification criteria*

Temperature criteria	
O4–B0	He I λ4471/He II λ4541 increase with type
	He I + He II λ4026/He II λ4200 increase with type
B0–A0	H lines increase with type
	He I lines reach maximum at B2
	Ca II K line becomes visible at B8
A0–F5	Ca II K/H_δ increase with type
	Neutral metals become stronger
	G band visible starting at F2
F5–K2	H lines decrease with type
	Neutral metals increase with type
	G band strengthens with type
K2–M5	G band changes appearance
	Ca I λ4226 increases rapidly with type
	TiO starts near M0 in dwarfs, K5 in giants
Luminosity criteria	
O9–A5	H and He lines weaken with increasing luminosity
	Fe II becomes prominent in A0–A5
F0–K0	Blend $\lambda\lambda$4172–9 increases with luminosity (early F)
	Sr II λ4077 increases with luminosity
	CN λ4200 increases with luminosity
K0–M6	CN λ4200 increases with luminosity
	Sr II increases with luminosity
Suffix notation	
e	Emission lines are present
f	He II λ4686 and/or CIII λ4650 in emission; mostly for O stars
k	Ca II K line when unexpected, e.g., interstellar in hot stars
m	Metallic; metal lines are stronger than normal
n	Nebulous; lines are broad and shallow; usually high rotation
nn	Very nebulous!
p	Peculiar; spectrum is abnormal
q	Queer; unusual emission; evolved from Q novae designation (archaic)
s	Sharp; lines are sharp, usually for early-type stars with low rotation
v	Variable; spectrum changes with time
w	Wolf–Rayet bands present (archaic)
Old prefix notation	
c for supergiants; *g* for giants; *d* for dwarfs	

classification is refined in successive steps. It is generally considered unwise to attempt an immediate accurate classification. Although key features are singled out, it is the appearance of the whole spectrum that is ultimately compared to the standards. Some care must be exercised, especially with early-type stars, because high rotation (see Chapter 18) can wash out weaker classification lines. Most modern spectral types are on the Morgan–Keenan system (MK) or a close derivative of it. See Johnson and Morgan (1953), Keenan (1963), and Gray and Corbally

(2009). A review of spectral classification is given by Morgan and Keenan (1973) and there is also a symposium dealing with this topic (Garrison 1984). Additional classification information can be found in the comprehensive objective-prism survey by Houk and Cowley (1975). Early O star classification is discussed by Walborn *et al.* (2002).

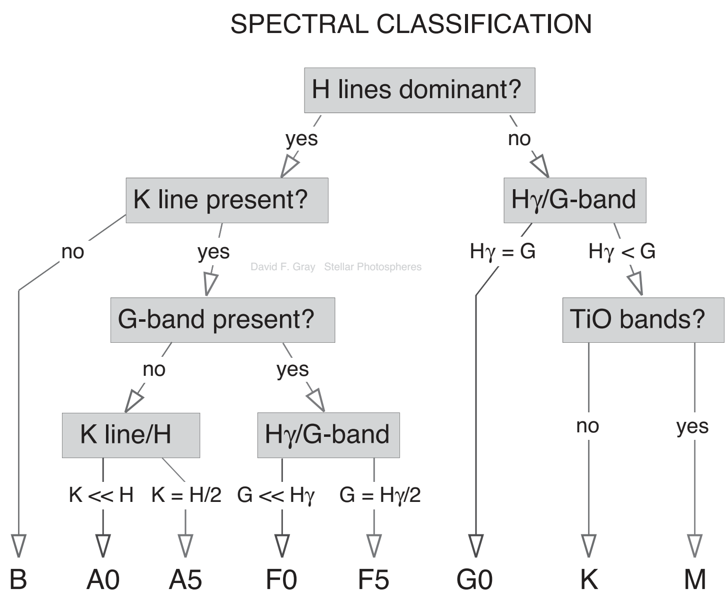

Figure 1.4 A flow chart organizes the classification basics. Start at the top.

Most catalogues of stars list spectral types, e.g., *The Bright Star Catalogue* (Hoffleit & Warren 1995), *Geneva Photometric Catalogue 1988* (Rufener 1988), or the SIMBAD database (simbad.u-strasbg.fr/simbad/).

Magnitudes and Color Indices

Stellar brightness is often expressed in magnitude units, defined by

$$m = -2.5 \log \int_0^\infty F_\lambda \, W(\lambda) \, d\lambda + \text{constant} \tag{1.4}$$

in which the amount of light or flux from the star, F_λ, is recorded in the spectral window specified by the response function $W(\lambda)$. The visual magnitude, often denoted by m_V or V, has its spectral window, $W_V(\lambda)$, centered on the yellow–green wavelengths near 5480 Å. Should we wish to include radiation at all wavelengths, i.e., $W(\lambda) = 1$ everywhere, the corresponding magnitude is called the *bolometric* magnitude. Even for this case, there remains an arbitrary zero-point constant. The constant is adjusted to suit an adopted set of standard stars (e.g., Landolt 2009, Smith *et al.* 2002), or, more recently, a well-calibrated standard star, usually Vega (α Lyr, HR 7001, A0 V), discussed in Chapter 10. In the latter case, all magnitudes in all color bands are set to zero for an A0 V star having $V = 0$. Vega itself has $V = 0.03$, so any magnitude defined for it is 0.03. Notice that the minus sign implies smaller magnitudes for brighter stars. The bright star Arcturus, for example, has $V = -0.05$, while the fainter star, σ Dra, has $V = 4.68$.

In many situations, there is no need to use the constant in Equation 1.4 because we deal with magnitude differences. Simplify the notation in Equation 1.4 and write

$$m = -2.5 \log F + \text{constant}. \tag{1.5}$$

That is, replace the integral by the flux, F, that is measured in the $W(\lambda)$ band. Now if two stars have magnitudes m_1 and m_2,

$$m_2 - m_1 = -2.5 \log \frac{F_2}{F_1}, \tag{1.6}$$

and the constant is gone. Equation 1.6 tells us that magnitude *differences* correspond to flux *ratios*.

Equation 1.6 can also be used on the same star in two different wavelength regions. A classic example is to measure the flux in a window in the blue, F_B, and compare it to the flux measured in the visible window, F_V. Then

$$m_B - m_V = B - V = -2.5 \log \frac{F_B}{F_V}. \tag{1.7}$$

The magnitude difference, $B - V$, is termed a color index. As you might have guessed, it is an empirical measure of a star's temperature.

Many standard magnitude systems have been set up. These differ in the wavelength positions and number of windows. One of the earliest is the red, green, ultraviolet or RGU system of Becker (1948). See also Becker and Stock (1954) and Becker 1963). The most commonly encountered systems are the UBV system (Johnson & Morgan 1953, Nicolet 1978) and its extension to longer wavelengths in the RIJKLMNH magnitudes (Johnson 1966, Johnson *et al*. 1966, Morel & Magnenat 1978, Landolt 1992), the uvby system (Strömgren 1966, Hauck & Mermilliod 1980, Perry *et al*. 1987), the Geneva $UBVB_1B_2V_1G$ system (Rufener 1988, Rufener & Nicolet 1988), and the DDO system (McClure & van den Bergh 1968). The UBVRI response functions are shown in Figure 1.5.

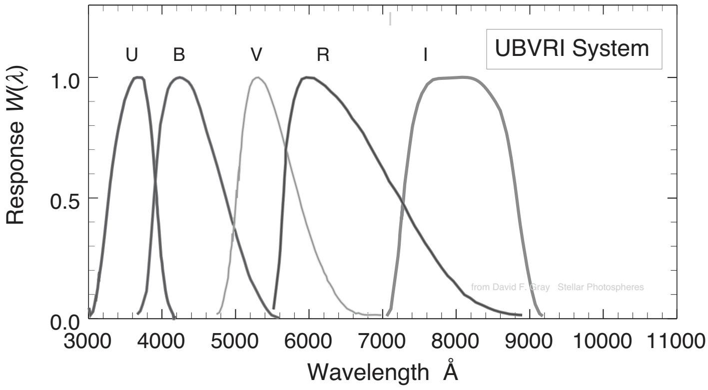

Figure 1.5 The UBVRI windows or response functions are shown normalized to unity. Based on Johnson (1966) and Bessell (2005).

More recently developed is the Sloan system (Fukugita *et al*. 1996, Ahumada *et al*. 2020). Its response functions are shown in Figure 1.6. There is less overlap between adjacent bandpasses compared to the UBVRI system. The Sloan observations extend to very faint magnitudes and cover hundreds of thousands of stars. Transformation equations between the UVBRI and Sloan systems are given by Krisciunas *et al*. (1998) and by Smith *et al*. (2002). A comprehensive review of photometric systems is given by Bessell (2005).

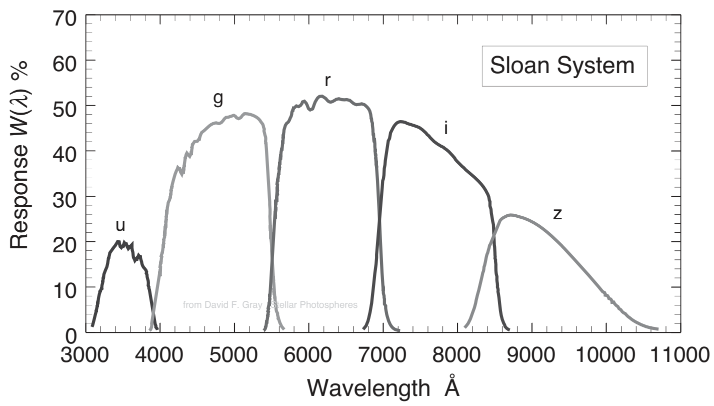

Figure 1.6 Response curves for the Sloan system are shown. These curves are the combined response of specially designed filters and the silicon response of a CCD light detector. Based on Fukugita *et al*. (1996).

Spectral types and color indices both give an empirical handle on stellar temperatures. Figure 1.7 shows the relationship between spectral type and the $B - V$ color index. Similar plots could be made for other color indices. A color index has the advantage of resolving finer temperature differences. The span from $B0$ to $M0$, for example has ~40 spectral-type classes, while the same span in $B - V$ can be divided into ~180 steps. Color indices can also be measured for considerably fainter stars, primarily a result of the ~1000 Å bands compared to the ~1–2 Å resolution elements in classification spectra. The primary disadvantage of color indices is interstellar extinction. Dust in space dims stars by progressively larger amount for shorter wavelengths. This makes stars seen through dust look redder or cooler. The extinction also makes stars look fainter. The spectral lines used for classification are not affected by extinction. The stars used for Figure 1.7 are nearby and have no interstellar reddening.

If both the color index and a spectral type are measured, we can go through Figure 1.7 and use the spectral type to predict the color index. Then we can compare the prediction with the measured value. The difference, observed minus predicted, is called the color excess, E_{B-V}, and it turns out to be related to the total extinction. In the *UBV* system, the V magnitude extinction, A_V, is

$$A_V = 3.1\, E_{B-V} = 3.1 \left[(B - V)_{\text{observed}} - (B - V)_{\text{predicted}} \right]. \tag{1.8}$$

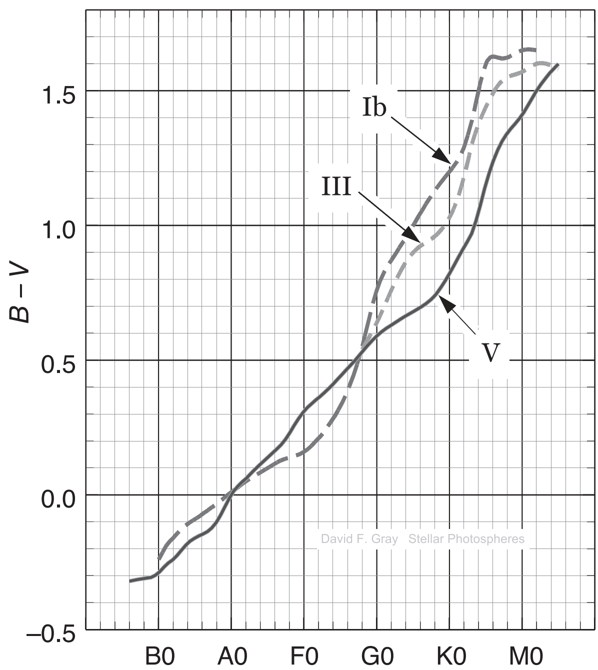

Figure 1.7 The mean relations between the $B - V$ color index and MK spectral types are shown based on the tabulation of Fitzgerald (1970).

The 3.1 coefficient in this relation is a nominal value (Hiltner & Johnson 1956, Whitford 1958, Schultz & Wiemer 1975, Krełowski & Papaj 1993). Larger values are found in some directions. For stars nearer than ~100 parsecs, there is little or no extinction.

There is one more important application for Equation 1.6 to which we now turn.

Stellar Distances and Absolute Magnitudes

How bright a star appears in the sky depends on how bright it really is and how far away it is from us. We can get the real relative brightness by mathematically allowing for the distance. The basic way to measure stellar distance is through the parallax introduced by the Earth orbiting the Sun. The effect is seen as an apparent displacement on the sky of nearby stars relative to stars much farther away. During the course of a year, nearby stars in the orbital plane of the Earth, the ecliptic plane, trace out a small straight line in the plane. Stars perpendicular to the plane at the north and south ecliptic poles trace out small circular motions (neglecting the small 0.017 eccentricity of the Earth's orbit). At positions in between, the apparent motion is an ellipse with its major axis parallel to the ecliptic, the so-called parallactic ellipse. Half the major axis of the ellipse is the parallax. From the star's perspective, this is the angular size of the Earth's orbit, or to be more precise, the angle subtended by one astronomical unit (a.u.) as seen from the star. The farther away a star is, the smaller its parallax. Call the parallax angle p arcseconds. The distance in units of