Chapter 13: **The Sun**

- Introduction / overview of the structure of the Sun
- Solar interior
- Photosphere
- Chromosphere and transition region
- The corona and magnetic activity; solar wind, heliosphere
- The solar activity cycle; magnetic activity on other stars
- sample problems

Intro / overview of solar structure

Our Sun is a middle-sized, middle-aged star. It has a hot core in which hydrogen fusion reactions produce helium and the energy that powers the Sun, and that powers the life on the surface of the Earth. Energy works its way slowly outward to the visible surface of the Sun, first by *radiation* (i.e., photons) and, nearer the surface, by *convection*, as buoyant bubbles of hot material rise. The visible surface, or *photosphere,* isn't the end of the Sun; the material above the photosphere is too low density to be easily visible but the Sun's *corona* extends many hundred of thousands of kilometers further. Streams of particles, the *solar wind*, flow outward past the planets and ultimately, at the edge of the *heliosphere*, collide with particles of the *interstellar medium*.

By mass the Sun started its life as ~71% H, ~27% He, and slightly less than 2% heavier elements (such as O, C, Ne, Fe). After ~4.6 billion years of fusion, the core composition has changed considerably, and is now estimated to be dominated by helium. The Sun is a hot fluid, hot enough that its atoms are mostly ionized; we call such an ionized fluid a *plasma*. The following table lists a few of the important properties of the Sun.

Table 13.1: selected properties of the Sun

Mass	$1.99 \cdot 10^{30}$ kg
Radius	$6.96 \cdot 10^{5}$ km
Luminosity	$3.828 \cdot 10^{26}$ J/s
Average density	1.41 g/cm^3
Temperature — Core	$15.7 \cdot 10^{6}$ K
Temperature — Photosphere	5,778 K ($T_{effective}$)
Temperature — Corona	range; several million K
Rotation period	~25 days near equator; ~35 days near poles
Magnetic field	~$1\text{-}2 \cdot 10^{-4}$ T globally; ~0.3 T in sunspots
absolute magnitude	$M_V = 4.83$
color index	B−V = 0.62
spectral type	G2V
metallicity	$Z = 0.0139$

The free electrons in the plasma interact with each other and with the magnetic field; one result is that the solar atmosphere is a strong source of radio emission, first detected in the early 1940s. The interior is opaque at radio wavelengths, just as it is at visible wavelengths; where, i.e., at what distance from the center, the opacity drops enough for photons to escape, is wavelength dependent. The photosphere, as noted above, is where the average visible photons can escape, and it's the photospheric temperature and radius that are meant, unless otherwise explicitly stated. In the radio, the Sun is hotter and larger: at ~1.4 GHz (21 cm) the majority of the radio emission escapes from the top of the chromosphere, where the temperature is nearing 10^5 K; at ~0.1 GHz (300 cm) the radio emission is predominantly coming from the corona, out where the temperature is ~2 million K.

The following sketch shows the structure of the Sun and the approximate conditions at the various boundary layers.

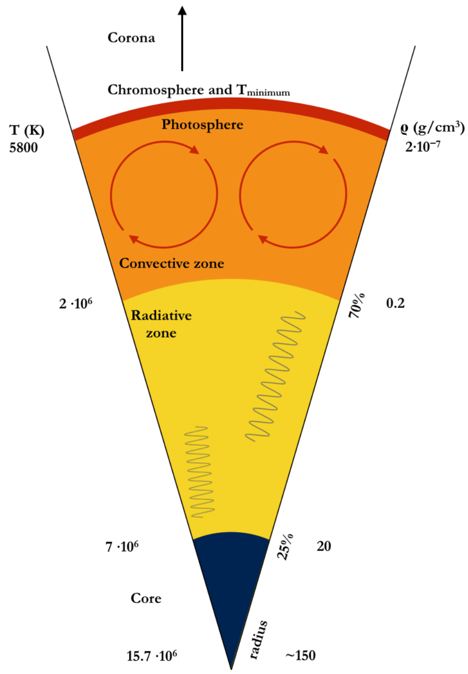


Figure 13.1: Solar structure

The solar interior

Energy is being produced in the core of the Sun predominantly by the proton-proton chain, described in the introductory chapter. Here are the main steps:

$^1H + {}^1H \rightarrow {}^2H + e^+ + \nu_e$

$^1H + {}^2H \rightarrow {}^3He + \gamma$

$^3He + {}^3He \rightarrow {}^4He + 2{}^1H$

We're using the notation for atoms, but remember that the core is hot and the atoms are all ionized. The two protons in the first step, which experience Coulomb repulsion because of their mutual positive electric charges, have to get close enough to each other for the strong nuclear force to fuse them into one nucleus (and then for the weak force to turn one proton into a neutron). It helps to be hot and dense, so that protons hit each other relatively often and relatively hard. As you may recall from the introductory chapter, the protons don't actually have to hit each other hard enough to overcome their Coulomb repulsion, just hard enough to get close enough for a quantum mechanical effect called tunneling, meaning that there is a significant probability that they are actually close enough to fuse. Turning to the electrons, if you are counting you can see that there are 4 e^- before (4 hydrogens) and 2 e^- after (1 helium). But each of the first steps in the proton-proton chain produces a positron (e^+), the electron's anti-matter twin. The first step happens twice in each ^{4}He fusion. The two positrons and the two spare electrons annihilate each other, producing a bit of the energy of the proton-proton chain.

On the order of 1% or so of the Sun's energy is generated via the CNO cycle, in which carbon serves as a substrate onto which hydrogens are added. The net result is the same, namely the fusion of $4{}^1H \rightarrow {}^4He$ + neutrinos and energy. The neutrinos leave immediately. Neutrinos travel at nearly the speed of light and interact very little with anything (more technically, neutrinos have a very low cross-section for interactions). It takes neutrinos only about 2 seconds to reach the solar surface.

Physics note: Neutrinos do interact via the weak nuclear force. By the late 1960s Ray Davis and John Bahcall had figured out that it should be possible to detect a small fraction of the solar neutrinos using chlorine. Davis installed a 100,000-gallon tank of perchloroethylene, C_2Cl_4, about 1500 meters underground in the Homestake Mine near Lead, South Dakota. Placing the detector deep underground reduces the likelihood of extraneous interactions with cosmic rays. A few energetic neutrinos will hit a chlorine atom, changing it into an argon atom: $\nu_e + {}^{37}\text{Cl} \rightarrow {}^{37}\text{Ar} + e^-$. The ^{37}Ar is radioactive, with a half-life of 35 days, and even though very few ^{37}Ar atoms were produced they could be counted. Bahcall estimated the number of detections Davis should make based on his best calculations of the rate of fusion reactions taking place in the Sun and the probability of a neutrino interacting with a chlorine atom (that very low cross-section). Davis' experiment detected only about 1/3 of the neutrinos Bahcall predicted. This was for many years known as the "solar neutrino problem", in part because one possible solution would have been that the Sun's fusion rate had somehow radically declined (and, because photons take a long time to work their way out of the Sun, we just hadn't noticed yet). The resolution came from advances in particle physics. Neutrinos come in three "flavors", displaying themselves as mu neutrinos and tau neutrinos, ν_μ and ν_τ, as well as the electron neutrinos to which Davis' experiment was sensitive. Later experiments were able to detect the other flavors. Our current understanding is that neutrinos "oscillate" among the three flavors. The Sun is far enough away and the oscillation rate rapid enough that the majority of electron neutrinos produced in the Sun will not arrive at Earth as ν_e, leading to the low numbers captured by Davis' early detector.

Unlike the neutrinos, no individual photon produced in the solar nucleosynthesis reactions is likely to make it to the surface of the Sun. If they did, we'd be seeing nothing but gamma rays. Rather, those photons are going to be scattered or they will be absorbed and re-emitted, often being directed back down inward, and slowly being stepped down in average energy. We say that the interior of the Sun is *opaque,* i.e., photons have a short mean free path between interactions. Inside the Sun conditions are very close to being a black body, meaning that if you could imagine being down inside the Sun the electromagnetic radiation that you'd see, that you'd be surrounded by, would have the appropriate black body spectrum for the temperature at your location. The temperature is gradually falling from the center to a minimum (~4,100 K) several hundred km above the level of the photosphere, so the photon energies must be falling as well. What starts out as gamma rays in the core will be predominantly visible frequencies by the time the energy reaches the photosphere.

There are three main methods energy can be transported from one location to another in a star: radiation, convection, and conduction. Radiation means photons. Convection refers to rising bubbles of material. Suppose that a bubble in the solar interior finds itself warmer than its surroundings. Being warmer, it will expand, and thus become less dense than its surroundings, and thus start to rise. As it rises, the surrounding pressure is falling, so the bubble expands further and as it expands it will cool. The temperature of the surrounding plasma has dropped also. The question for convection is whether the temperature in the bubble remains higher than the temperature of its surroundings. If the surroundings cool more rapidly than the bubble, the bubble will remain buoyant and continue to rise. In the Sun conditions are right for convection in the outer ~30%, by radius, of the interior. The inner ~70% is dominated by radiative energy transport. (When we consider stars more generally we'll return to convective zones and see that less massive stars have relatively deeper convective zones and slightly more massive stars have thinner ones; very massive stars have convective cores rather than convective outer envelopes.) The third energy transport mechanism, conduction, occurs when hot, jiggling particles bump into other nearby particles and transfer some of their kinetic energy in the process. In stars conduction is relevant for stellar remnants such as white dwarfs, but not for the solar interior.

The interior of the Sun is dense enough for acoustic waves to propagate. The Sun rings like a bell, in many thousands of frequencies all at once. Just as we can use seismic waves to probe interior conditions on Earth, the study of *helioseismology* helps probe interior conditions in the Sun. Particularly apparent to us are the 5-minute oscillations, pressure waves with frequencies in the 2-4 mHz regime. The motions and amplitudes of these waves give us insights into the sound speed, and hence the density, at various depths in the Sun. Analysis of the waves suggests that the radiative zone rotates uniformly, as if it were a solid, unlike the convective zone which exhibits *differential rotation*, rotating more rapidly at lower latitudes. The transition layer between the convective and radiative zones is called the *tachocline*. The tachocline at the base of the convective zone is thought to be the principal location for the generation of the Sun's magnetic field. Recall here that moving electric charges, which we have in the rotating plasma in the Sun, produce a magnetic field. We'll examine the visible effects of the Sun's magnetic field when we consider the corona and the solar activity cycle, below.

The photosphere

At the top of the convective zone is the thin layer called the photosphere (scale height ~150 km), because this is the layer from which the average photon stands a better than average chance of escaping. The density has dropped to the point where rising convective bubbles are not longer supported. We see the bright tops of the bubbles in a pattern called *granulation*; hot material burbles up, slides sideways, cools, and sinks back down into the interior. Granules are approximately 1,500 kilometers across and individual granules last for about 15 minutes. Some of the most detailed images of the photosphere have been acquired with the Swedish 1-m Solar Telescope on La Palma, in the Canary Islands. The following image was taken in August 2003; the bar across above the sunspot is roughly twice the diameter of the Earth. The granules are surrounded by darker, cooler lanes of sinking plasma. There are also some enigmatic bright points around the edges of the granules as well, where the magnetic field seems to be stronger than average, and where we are seeing deeper than average into the Sun.

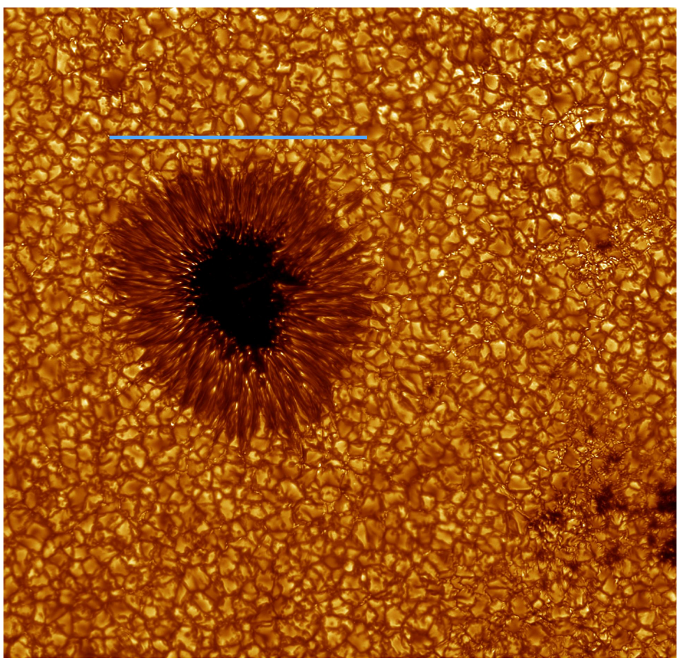


Figure 13.2: Solar granulation and sunspot.

Credit: G. Scharmer, K. Langhans, M. Löfdahl, Institute for Solar Physics, Swedish Solar Telescope image. AR 425 / 8 Aug. 2003 / 436.4 nm http://www.isf.astro.su.se

Because the photosphere isn't a solid surface we see deeper into the Sun when we look at the center of the Sun's disk and less deep when we look toward the limb. An uncorrected image of the full disk of the Sun will exhibit limb darkening. The following sketch illustrates this.

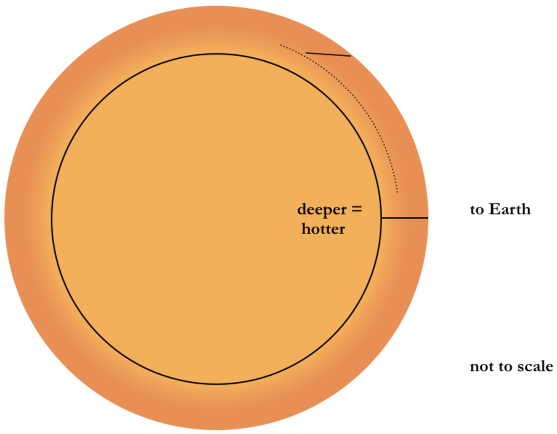
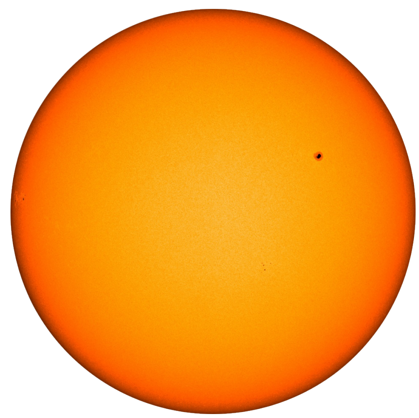


Figure 13.3: Limb darkening. At disk center we see light that comes from deeper, hotter layers than the light we receive from near the limb.

Figure 13.4: The image at right, from the Solar Dynamics Observatory, 31 January 2016, shows the limb darkening effect.

Credit: NASA / SDO and the HMI science team.
http://sdo.gsfc.nasa.gov/data/



How deep into the Sun we can see is dependent on wavelength. At a wavelength that falls in the middle of a strong absorption line we aren't seeing very deep. For instance, consider two of the strongest absorption lines in the Sun, the H and K lines of singly ionized calcium, at 397 and 393 nm, respectively.

Historical / terminology note: By the early 1800s observing technology had advanced to the point at which it became possible to detect the strongest of the absorption lines in the solar spectrum. Fraunhofer, one of the early solar spectroscopists, measured and labeled several hundred of the solar absorption lines. The most prominent lines he designated with capital letters, A through K (with no I or J), starting at the red end of the spectrum. Over the decades various investigators determined which elements were the source(s) of which lines. Some of Fraunhofer's terminology is still in use — the D lines, for instance, are sodium lines, G is a tight band of lines including some due to the molecule CH, and H and K, as noted above, are lines of singly ionized calcium; on the other hand, people rarely use Fraunhofer's C or F when they are referring to Hα or Hβ. Some of his lines turned out to be telluric lines, i.e., actually due to absorption by oxygen molecules in our atmosphere, not due to the Sun.

Another terminology note: In addition to using superscripts to indicate ions — e.g., H^+ — it is also standard spectroscopic practice to denote ionization state with Roman numerals. In this notation, neutral hydrogen would be H I and ionized hydrogen would be H II. Using this shorthand we can write Ca II K to denote that 393 nm line of singly ionized calcium.

To get back to the H and K lines, there is a high probability that photons at 397 and 393 nm will interact with a calcium ion; the *opacity* at those wavelengths is relatively high and the *mean free path* of a photon at those wavelengths is relatively low. By the time a photon at these wavelengths stands an even chance of escaping it's out around the temperature minimum, several hundred kilometers above the average level of the photosphere. In the nearby continuum part of the spectrum, i.e., at wavelengths not within these absorption lines, the opacity is much lower, photons have longer mean free paths, and they are likely to escape from lower down in the photosphere where the temperature is higher. In other words, we are seeing deeper into the Sun in the continuum part of the spectrum than we are in the lines.

There are also some significant sources of continuous opacity, meaning absorption that takes place over a large range of wavelengths and not just in an absorption line. A source of continuous opacity that's very important in the solar photosphere is ionization of the H^- ion. It is possible to form a negative hydrogen ion, although that second electron is not very tightly bound — the ionization energy is only ~0.75 eV. Ionization of H^- is a major source of opacity in the visible and near infrared portions of the solar spectrum. Having stressed the fact, above, that the interior of the Sun is a plasma, it's worth noting that in the photosphere the vast majority of the hydrogen atoms are neutral rather than positive ions. The more easily ionized elements, such as Fe or Ca, are still more likely to be ionized (and to supply electrons to make H^-), but recall that it takes 13.6 eV to remove that ordinary electron from hydrogen. Conditions are just not hot enough or energetic enough for most of the hydrogen to be ionized in the solar photosphere (which is why it's possible for there to be negative hydrogen ions).

Given that much of the hydrogen is neutral and that the Sun has plenty of light in the visible part of the spectrum, there is a temptation to expect strong hydrogen Balmer absorption lines. Resist that temptation! Recall that the Balmer lines are absorptions that arise when the hydrogen electron is *already* in the second energy level and that it takes an ultraviolet energy (~10.2 eV) to excite the electron from the ground state into that second energy level. The solar photosphere isn't hot enough — there are not enough UV photons — for a high percentage of the hydrogen atoms to have their electrons in the second level. Here is an image of the visible spectrum of the Sun; the dark line 7 rows down is Hα; the two strong lines in the yellow are due to sodium; many of the lines in the green and blue are due to Mg and Fe (and Hβ and Hγ); the Ca II H and K lines are just too short to show in this image:

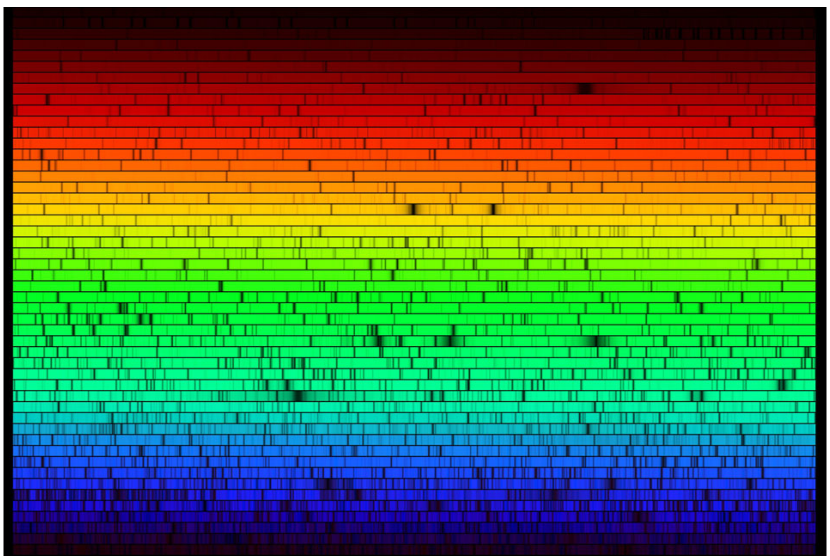


Figure 13.5: The solar spectrum, obtained at the McMath-Pierce Telescope at the National Solar Observatory at Kitt Peak. Credit: Nigel Sharp / NSF / NSO / AURA.

https://www.noao.edu/image_gallery/html/im0600.html

The following sketch shows an average temperature structure for the atmosphere of the quiet Sun, meaning not above sunspots or other regions that show strong magnetic activity. It starts at the depth at which the atmosphere becomes opaque at 500 nm. It's based on work done by Eugene Avrett, of the Smithsonian Astrophysical Observatory, and colleagues. The base of the photosphere is usually taken to refer to the top of the convective zone. When we talk about the temperature of the photosphere we usually mean the temperature at the base, not the temperature minimum where the photosphere fades into the chromosphere.

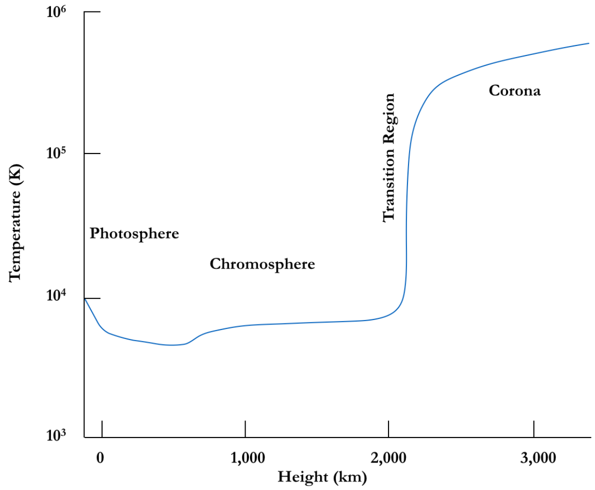


Figure 13.6: Temperature structure of the solar atmosphere.

Chromosphere and transition region.

The temperature minimum is roughly 500 km above the base of the photosphere and the chromosphere extends for over a thousand kilometers above that. In the chromosphere the temperature gradually rises again to roughly 10^4 K, until we reach the transition region, where the temperature increases rapidly to the millions of degrees which characterize the corona. The temperatures (and the low density) in the chromosphere favor emission in several strong lines, particularly the red Balmer Hα line. Arches of bright red hydrogen visible during total solar eclipses are the reason for the name chromosphere — *chromo* means color. Observing the disk of the Sun in white light, the light from the chromosphere is overwhelmed by the very much brighter underlying photosphere.

If we observe with narrow bandpass filters so that we can isolate specific emission lines, such as Hα or the Ca II H and K lines, it's possible to take pictures of the solar chromosphere. The chromosphere is not a nicely defined layer of distinct thickness; spicules (spikes a few hundred kilometers high) and arches of chromospheric material of held aloft by magnetic fields (prominences or filaments) stick up into the lower corona. The following image, taken in an emission line of ionized helium (He II 304 Å) by the space-based Solar Dynamics Observatory shows some of the structure of the upper chromosphere.

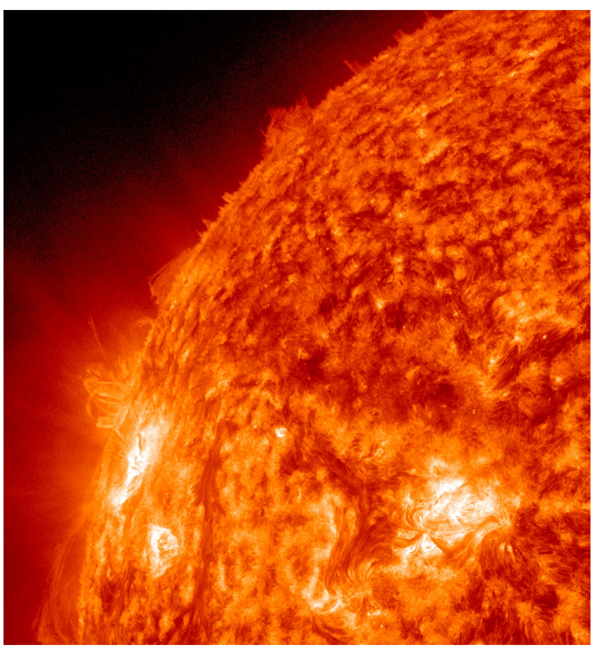


Figure 13.7:
The Sun in He II 304 Å 11 November 2012. Courtesy of NASA / SDO and the AIA science team.

http://sdo.gsfc.nasa.gov

The spicules, particularly visible along the limb in the upper part of this image, seem to outline the edges of supergranules, convection patterns that are similar to granules but larger, roughly 25-35,000 km across, with somewhat lower vertical velocities. Around the edges, where material is sinking, the magnetic field also collects and gets enhanced and supports the chromospheric plasma we see, particularly at the limb, as spicules.

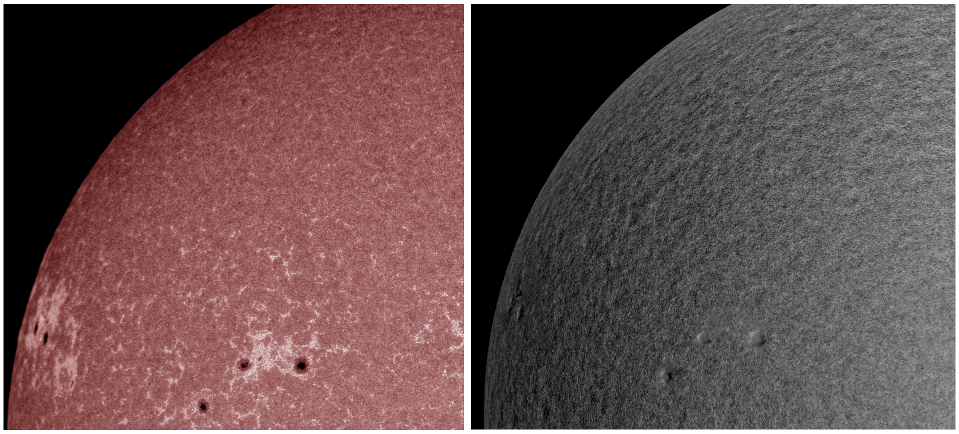


Figure 13.8: The Sun a) at 1700 Å (continuum) and b) as a Dopplergram, 11 November 2012. Courtesy of NASA / SDO and the AIA and HMI science teams. http://sdo.gsfc.nasa.gov

The two images above are of the same portion of the Sun and taken at roughly the same time as the He II 304 Å image. The left-hand image is representative of the temperature minimum; you can see dark sunspots, nearby less well-defined bright areas called *plage*, and the generally mottled appearance of the surrounding quiet Sun. The right-hand image is a Dopplergram and shows, using differing shades of gray, where material is rising and where it is falling. We'll look at sunspots in more detail when we consider the Sun's magnetic activity, but note here that their stronger magnetic fields act as plugs, inhibiting convective motion.

On a related note, if we look at the whole Sun and display the Doppler shifts using these shades of gray you can see the solar rotation as well as that mottling due to the rising and falling regions of the surface. Here's a full-disk Dopplergram for 11 November 2012:

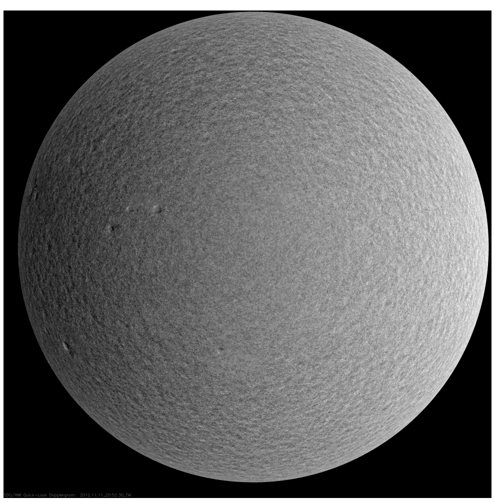


Figure 13.9: Solar Dopplergram, 11 November 2012. Courtesy of NASA / SDO and the HMI science team.

http://sdo.gsfc.nasa.gov

The arches of chromospheric material held up by the magnetic field are called *prominences* when we see them above the limb of the Sun. They are the same structures, looking relatively bright when seen against the black of space above the limb of the Sun, while looking relatively dark when seen against the hotter underlying photosphere (where they are called filaments). The following He II 304 Å image is just south of the region show in the images above. A bright arch of material rises above the limb and a darker filament can be seen against the disk of the Sun. Prominences can dwarf the Earth in size and last for weeks. Short spiky spicules are visible along the bottom of the Sun's disk; they last for ~15 minutes.

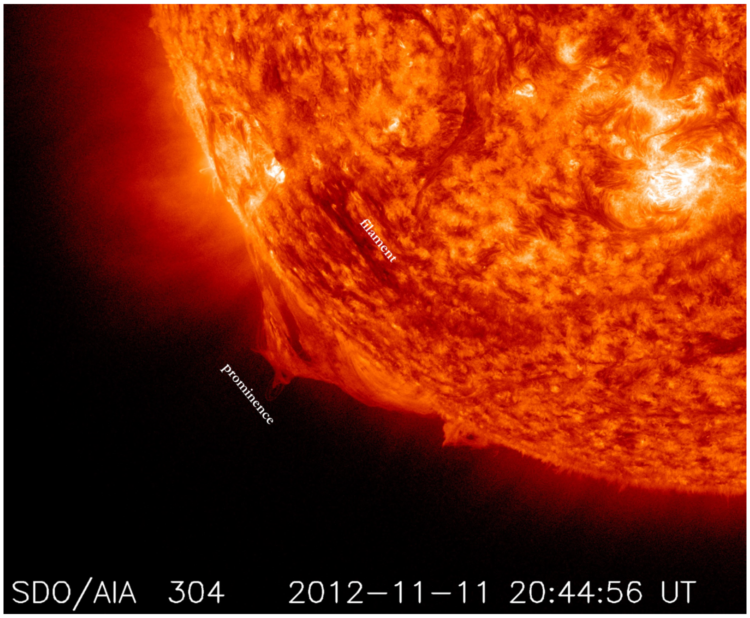


Figure 13.10:
The Sun in He II 304 Å 11 November 2012. Courtesy of NASA / SDO and the AIA science team.
http://sdo.gsfc.nasa.gov

The corona and magnetic activity; solar wind

In the upper chromosphere the temperature starts to rise, soon reaching tens of thousands of degrees. Then, within a few hundred kilometers, the temperature soars up over a million Kelvin and we are in the corona. The corona is very hot but its density is even lower than the chromospheric density (the corona would be a really good vacuum on Earth), tailing off at a distance of several solar radii into the *solar wind*, a somewhat steady / sometimes clumpy outflow of particles from the Sun. Near Earth we measure the density of the solar wind to be a few protons / cm^3. More on the solar wind, below. Like the chromosphere, the corona is visible during solar eclipses or in emission lines. In this case, the emission lines are mostly due to highly ionized species. It's also easier to image the corona in the far UV or X-rays because at those short wavelengths the underlying photosphere is not so bright; of course, we need to be in space to do this. In the visible we tend to see the corona either by photospheric light being scattered by fast-moving free electrons or by dust. The light scattered by the electrons is called the K-corona, from the German term for continuous; the spectrum is continuous because the electrons are moving so rapidly that any photospheric spectral features are smoothed out by Doppler effects. The light reflected off the dust, which dominates the visible corona farther from the Sun, will show the photospheric lines and so is called the F-corona, for Fraunhofer.

The Solar and Heliospheric Observatory, SOHO, launched in December 1995, has been a reliable source of solar coronal images, still functioning well in 2019. SOHO is a joint mission of NASA and ESA (the European Space Agency). SOHO is equipped with a *coronagraph*, an instrument that blocks out the light from the disk of the Sun to produce artificial eclipses. The following two images show what the corona, to two different distances, looked like on a few hours after the November 2012 time of the above images.

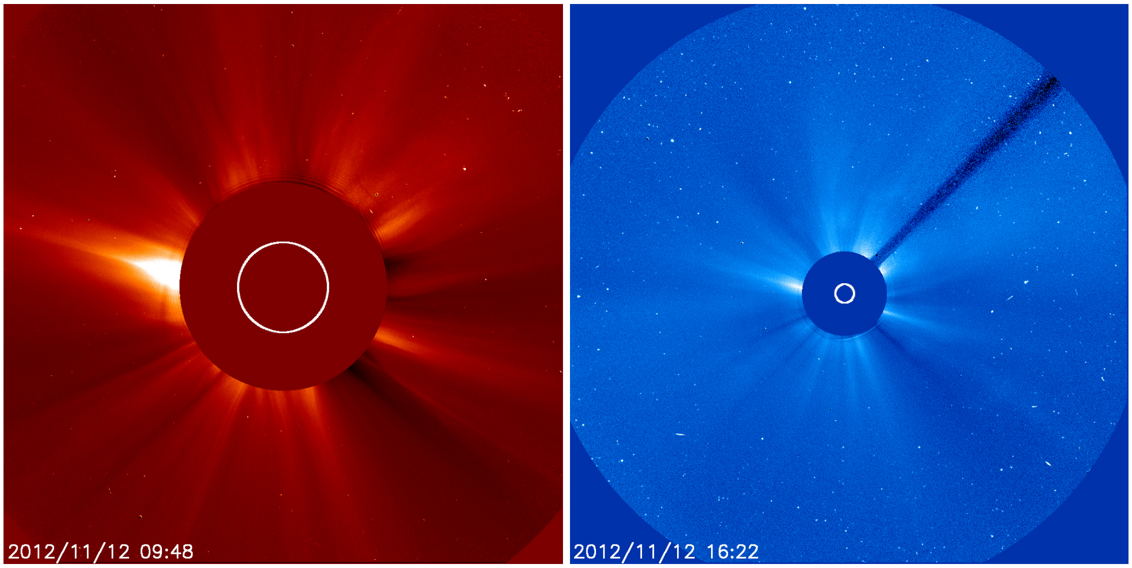


Figure 13.11: Solar corona; a) inner C2 image, b) wider field C3 image. Credit: Courtesy of SOHO / LASCO consortium. SOHO is a project of international cooperation between ESA and NASA. http://sohodata.nascom.nasa.gov/cgi-bin/data_query

In both images the visible disk of the Sun is indicated by the white circle. The left-hand, C2 image, shows the corona out to ~8.4 $\cdot 10^6$ km; the C3 field of view, on the right, is ~45 $\cdot 10^6$ km, or about 32 solar diameters across. The dark line in the C3 image is the support holding the disk of the coronagraph. The structure of the corona is governed by the Sun's magnetic field and varies over the course of the solar cycle. The streamers, which are associated with magnetic active regions, are concentrated nearer the Sun's equator at solar minimum and are also present at higher latitudes at the peak of the activity cycle. When the Sun is less active we see *coronal holes,* usually, but not always, near the poles.

Seeing the corona off the limb of the Sun has always been possible — just wait for a total solar eclipse. But imaging the corona in front of the disk of the Sun isn't easy. The corona is hot, many atoms are highly ionized, and the emission from those atoms will be in the extreme ultraviolet. One benefit of this is that the underlying photosphere isn't very bright at such short wavelengths so there isn't much competition from the photosphere; the main downside is that the photons we want to capture are so energetic. Arthur Walker Jr. (1936 - 2001), an African-American solar physicist / optical engineer, developed the technology to obtain full-disk images of the Sun in the extreme UV using normal-incidence optics. Walker launched his cameras on sounding rockets, i.e., short, non-orbital flights, and was able to capture the first high-resolution full-disk images of the solar corona.

Today some of the best ultraviolet images of the inner corona are provided by the Solar Dynamics Observatory, launched by NASA in February 2010. The following images compare the corona on January 1st, 2016 with the corona on November 11th, 2012. These images are taken at 211 Å, light emitted from Fe XIV at ~ 2 million K. In 2016, the coronal holes are much more extensive.

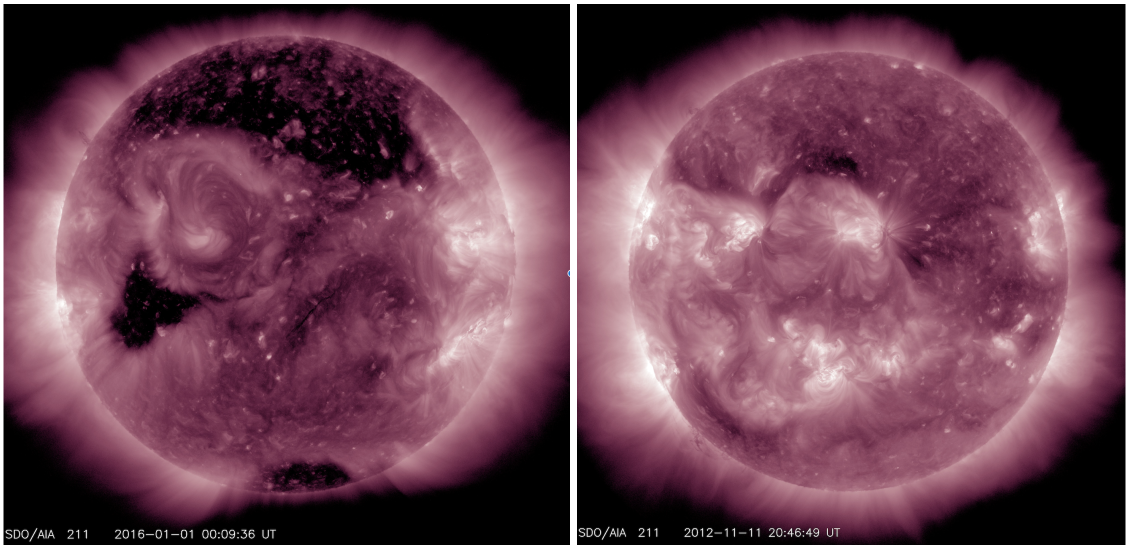


Figure 13.12: Coronal Holes; 211 Å, a) 1 January 2016 and b) 11 November 2012. Courtesy of NASA / SDO and the AIA science team. http://sdo.gsfc.nasa.gov

The following two images are again of the Sun on November 11, 2012. The left-hand image is Fe IX 171Å, most strongly emitted at temperatures of $\sim 6 \cdot 10^5 - 1 \cdot 10^6$ K, roughly 5,000 km above the temperature minimum. The right-hand image is a magnetogram, which uses shades of gray to represent the strength of the magnetic field. You can readily see that the regions that are bright in the corona correspond to regions of stronger magnetic fields, and, if you refer to the continuum image above, to the locations of sunspots in the photosphere.

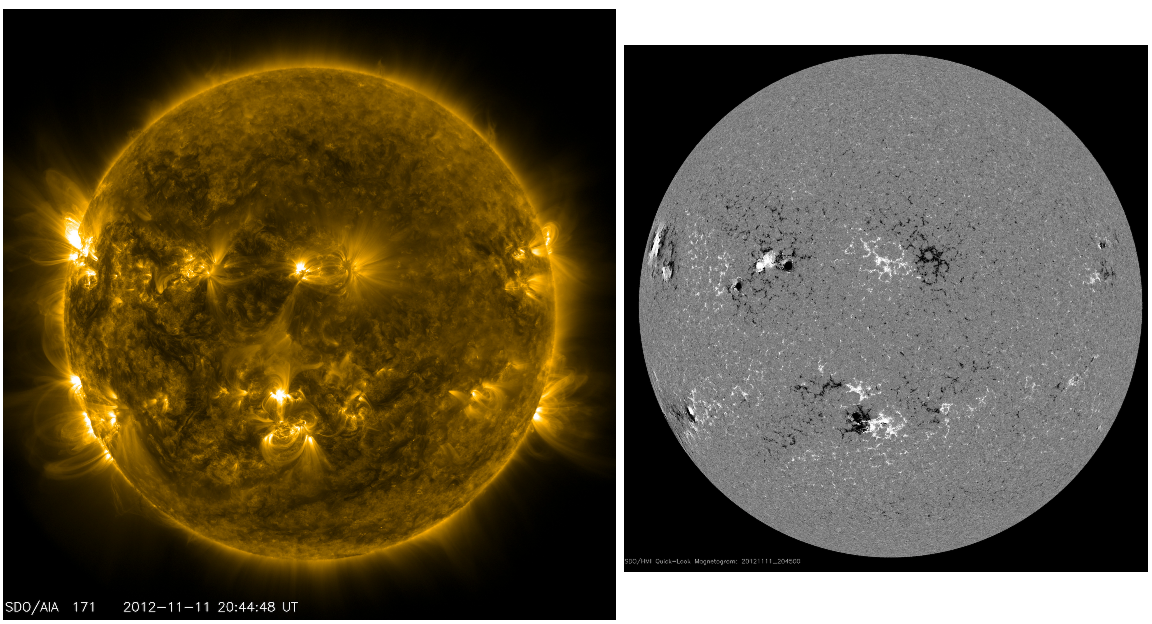


Figure 13.13: The Sun a) at 171 Å and b) as a magnetogram, 11 November 2012. Courtesy of NASA / SDO and the AIA and HMI science teams. http://sdo.gsfc.nasa.gov

Let's look in detail, above, at the south central active region in the above Fe IX 171Å image. You can see the loops in the emitting plasma; it's being supported by the magnetic fields. Solar physicists with the SDO model the structure of the Sun's magnetic fields. The second image, below, shows what the field looked like around that active region.

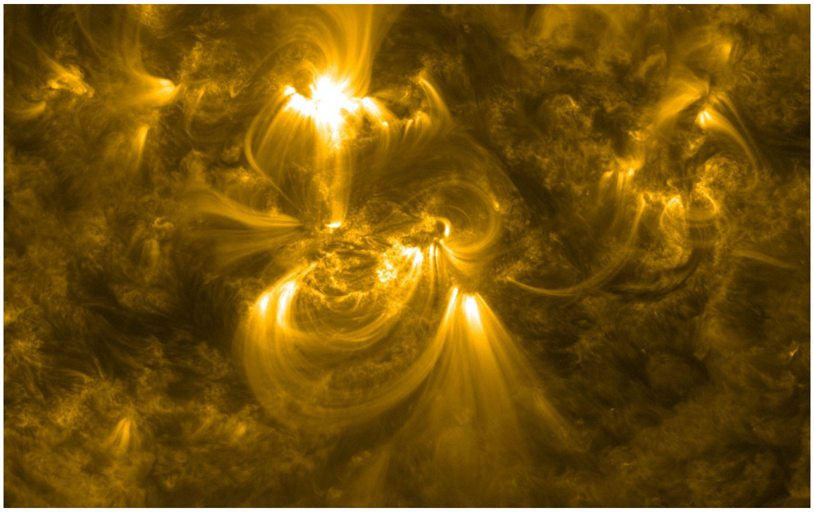


Figure 13.14a: Active region on the Sun at 171 Å 11 November 2012. Courtesy of NASA / SDO and the AIA science team. http://sdo.gsfc.nasa.gov

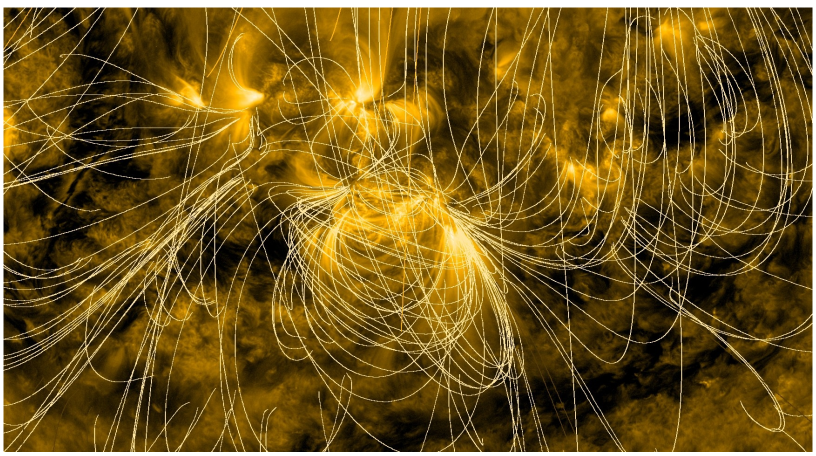


Figure 13.14b: Active region and magnetic field lines on the Sun at 171 Å on 11 November 2012. Courtesy of NASA / SDO and the AIA and HMI science teams. http://sdo.gsfc.nasa.gov

We have learned in recent decades that the magnetic field is intimately connected both with the structure and the heating of the solar corona. There are a few emission lines in the visible that are indicative of the high level of excitation of the corona and that can be observed from the ground during solar eclipses or using a coronagraph. One of the first to be identified is often just called " the green line", at 5303 Å. Nobody was expecting such high levels of ionization and it wasn't until the early 1940s that this emission line was identified as being due to Fe XIV (before that the unknown element responsible was simply called "coronium"). After all, as soon as it was clear that the Sun produces its energy in the core, it made total sense that the temperature would fall with increasing distance from the center.

Physics note: An additional reason the 5303 Å line was tough to identify is that it is a forbidden transition, meaning that this spontaneous emission violates quantum mechanical selection rules. Most excitations are followed almost immediately (e.g., 10^{-8} s) by de-excitations, but some energy levels, called metastable levels, don't have a permitted de-excitation route. An electron may have a lifetime in a metastable level many orders of magnitude longer than in a level from which spontaneous emission is permitted. In the relatively high-density laboratory setting atoms are jostled around enough that electrons are usually excited up out of a metastable level long before they might drop downward and emit so we have less experience observing these transitions. In the vacuum of the corona, de-excitation is possible.

It was, as noted, a bit of a shock to realize that in the corona the temperatures go back *up* to temperatures that, at least in some places, are as high as the temperatures in the center of the Sun. The low densities mean that once an atom becomes highly ionized, it's less likely to run into an electron and recombine; collisions are also important in cooling, since energy can be released in a collision, so lower densities and fewer collisions make it easier to maintain high coronal temperatures. But how the corona gets so hot in the first place isn't obvious. The lower-energy visible wavelength photons from the photosphere won't do it. Neither does it seem reasonable to heat the corona mechanically; e.g., it would see that the corona would need to be denser for it to be possible for acoustic waves associated with the convection to carry energy into the corona. Some recent studies have suggested that, despite the low density, some modes of the solar oscillations may well propagate into the corona where, with the change in density, they become shock waves capable of dissipating heat. Recent observations also suggest that there is a continual low level of flaring, in which energy is released from the magnetic field. Magnetic field lines snap and reconnect in different, lower-energy configurations, releasing energy that heats the surrounding plasma. The Parker Solar Probe (PSP), named for solar physicist Eugene Parker (d. 2022), who first proposed the existence of the solar wind, was launched in August 2018. It was placed into a close, eccentric, and, with some help from Venus flybys, shrinking, orbit around the Sun. At its perhelion passage in June 2025 it passed about 6.2 million km of the Sun's surface, deep enough into the corona to provide valuable data about why the coronal temperature is millions of kelvin. Initial results suggest that Alfvén waves play a role in coronal heating. Alfvén waves are plasma, or magnetohydrodynamic, waves involving the oscillation of ions; the restoring force is provided by a tension in the magnetic field. The name dates from a 1942 suggestion by plasma physicist Hannes Alfvén that it should be possible for a plasma wave to carry energy from the photosphere out into the corona. The PSP is expected to remain in orbit around the Sun and to carry out observations of the solar atmosphere as we enter the declining phase of the current sunspot cycle.

More information about the corona and solar wind will come from the Solar Orbiter, a primarily ESA (European Space Agency) probe launched in February 2020. In June 2020 its initial perihelion passage took it about 0.5 AU from the Sun, where it captured images of hot spots nicknamed "camp fires" that may be involved in coronal heating. Like the Parker Probe, Solar Orbiter will also use gravitational interactions with Venus to modify its trajectory, making its first flyby of Venus in December 2020.

Let's look again at the Sun at 171 Å on November 11, 2012. In the following image you can see material in bright loops near the limb. The cartoon on the right shows what may be happening in magnetic reconnection.

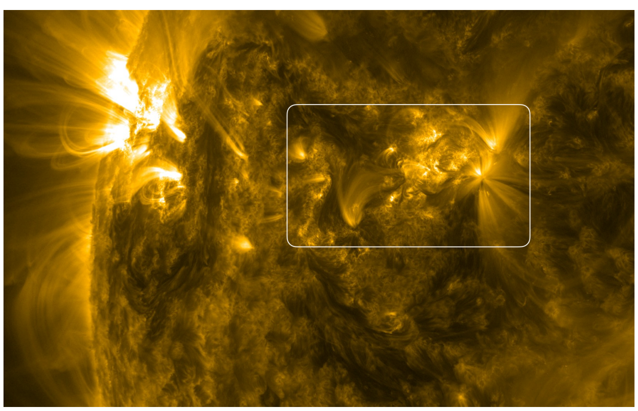


Figure 13.15: The Sun at 171 Å, 11 November 2012. Courtesy of NASA / SDO and the AIA science team. http://sdo.gsfc.nasa.gov

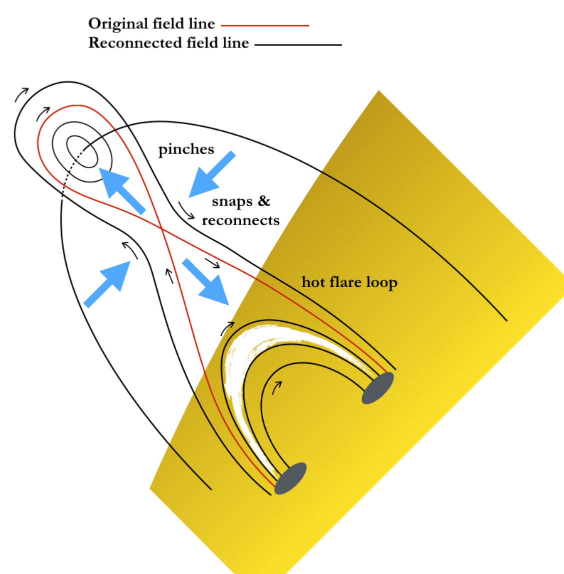


Figure 13.16

Flaring, brightenings that occur over a few minutes, happens at smaller scales, as well. In the next two images, we'll focus on the region highlighted by the box, above. Look at the regions that brighten from one frame to the next, taken 15 minutes later. Whether this type of heating is adequate to explain the coronal energy budget is still being debated.

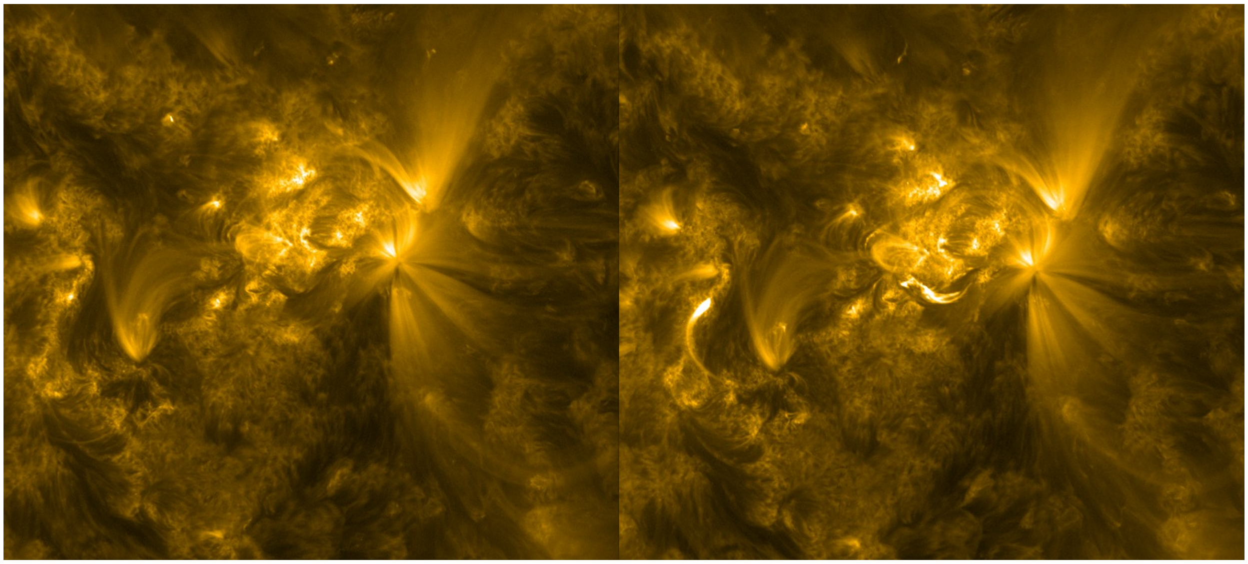


Figure 13.17: Small-scale flaring. The Sun at 171 Å, 11 November 2012, a) 23:44 UT and b) 23:59 UT. Courtesy of NASA / SDO and the AIA science team. http://sdo.gsfc.nasa.gov

Both flares and prominences can be associated with *coronal mass ejections*, which, as the name suggest, involve substantial amounts of plasma lifting off into space. In a flare, for instance, the material above the reconnection point may by ejected. When a prominence erupts, it's often the case that some of the prominence material also drains back down into the lower atmosphere. A filament erupted in spectacular fashion on August 31, 2012, when the region just beneath it flared. The following set of three images are in He II 304 Å, spaced about 15 minutes apart.

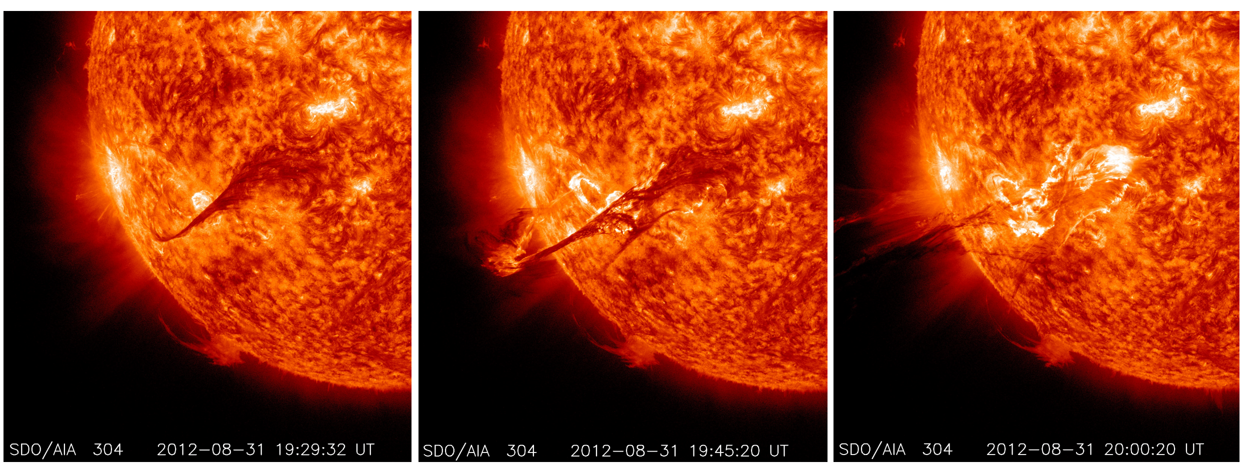


Figure 13.18: Coronal Mass Ejection. The Sun at 304 Å, 31 August 2012; a) 19:29 UT, b) 19:45 UT, c 20:00 UT. Courtesy of NASA / SDO and the AIA science team. http://sdo.gsfc.nasa.gov

Coronal mass ejections inject huge amounts of plasma into the *solar wind*, the otherwise fairly steady outflow of charged particles from the Sun. In the late 19th and early 20th centuries three types of observations in particular pointed to the existence of the solar wind. The first is that the corona could be observed during solar eclipses to extend many solar radii out into space, with no clear end. A second observation is that the gas tails of comets point away from the Sun. Thirdly, flares on the Sun are often followed, a bit more than a day later, by stronger than average aurorae and other evidence of geomagnetic disturbances. In the aurorae, charged particles from the Sun are deflected by the Earth's magnetic field, slamming into our atmosphere near the north and south magnetic poles where they excite air molecules. Oxygen and nitrogen emit several shades of green, red, and violet as they de-excite. This image was taken from above the aurorae.

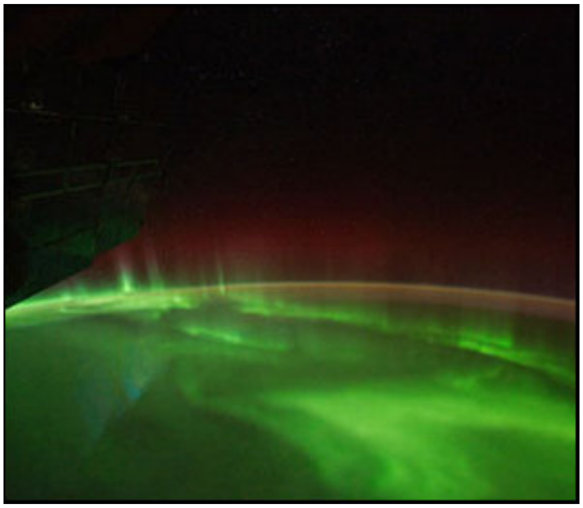


Figure 13.19: The Aurora Australis from the International Space Station. NASA, 2011. http://spaceflight.nasa.gov/gallery/images/station/crew-29/html/iss029e007502.html

The Sun loses on the order of 10^{-14} $M_{\odot}$ per year; some stars have much stronger stellar winds with much higher mass loss rates. In part because it is so tenuous coronal material doesn't cool easily, meaning that it stays hot and particles continue to have high speeds even far out in the corona where the Sun's gravitational tug has lessened considerably. Thermal speeds and energy from the magnetic fields both seem to play a role in accelerating some coronal particles to escape speed, feeding the solar wind.

Particles from coronal streamers tend to be slower than particles that leave through coronal holes (~400 km/s vs. ~750 km/s). Satellite data confirm that the slow solar wind particles have a composition that matches the corona and a temperature of ~ $1.5 \cdot 10^6$ K. As noted above, near solar minimum the streamers tend to be confined to latitudes closer to the equator (within about 30 degrees or so) and are seen much more evenly distributed in latitude at solar maximum. The fast solar wind more closely matches the upper photosphere in composition, and is cooler and less dense than the slow solar wind. Near the Earth the density of the solar wind is usually a few protons per cm^3.

The Sun (as well as other stars with out-flowing plasma winds) is surrounded by a bubble called the *heliosphere* (or astrosphere for other stars), within which the pressure of the of the solar wind exceeds that of the gas of the interstellar medium. The extent of the heliosphere varies depending on the strength of the solar wind and the density of the local interstellar medium; its radius is roughly 120 AU, meaning that the major planets lie inside the heliosphere but that some Trans-Neptunian objects with more eccentric orbits will travel beyond it — e.g., Sedna, with an aphelion of roughly 900 AU. The Oort cloud, the proposed cloud of icy planetesimals and source of long-period comets, lies well beyond the heliosphere's boundaries. Particles in the solar wind encounter ISM particles and slow relatively abruptly at a layer called the termination shock, the inner edge of a turbulent region called the heliosheath. At the outermost layer of the heliosphere, called the heliopause, the pressures of solar wind and ISM particles have finally equalized. Voyager 1 appears to have crossed the heliopause in August, 2012, and entered interstellar space at a distance of 121 AU from the Sun. Voyager 1 is still going, over 142 AU from the Sun by June, 2018, and, more than 40 years since its launch, still communicating with us. Voyager 2 is on a trajectory that's taking it out of the solar system a bit more slowly than its twin (3.3 AU/year, compared to Voyager 1's 3.6 AU/yr) and is expected to reach the heliopause in early 2020. Both spacecraft are expected to have enough power to continue to transmit some data about their environments until the mid-2020s.

The solar activity cycle

Individual sunspots last for a few days or a few weeks. Spots are caused by the Sun's magnetic field. When it is at its most symmetric the Sun's magnetic field resembles the dipole field of a bar magnet, shown below.

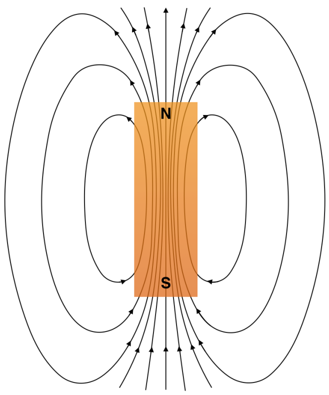


Figure 13.20: The magnetic field lines in and around a bar magnet.

In the Sun, though, we have a differentially rotating, convecting plasma. In the solar interior the plasma is dense enough to drag the magnetic field lines around, introducing helicity to the field lines and generating a toroidal component to the magentic field. Over time, the field lines get wrapped around the Sun, twisted and kinked. Loops of magnetic field poke up through the photosphere. Where the magnetic flux is concentrated, convection is inhibited. At the tops of convective bubbles plasma rises, moves sideways, and sinks. A bundle of magnetic field lines sticking up perpendicular to the level of the photosphere inhibits that sideways motion of the plasma. As a result, the material in a sunspot cools off and stays put, preventing hot material from rising to the surface from deeper in the convective zone. Sunspots are approximately 1,500 - 3,000 K cooler than the surrounding quiet Sun.

Sunspots start as *pores*, on the order of a few tens of kilometers across. Large spots can be many times the diameter of the Earth. A spot has a central dark *umbra*, surrounded by a wispy region called the *penumbra*. Spots often come in pairs, not necessarily of equal size, with opposite magnetic polarities, a bit as if a horseshoe-shaped magnet were sticking up out of the photosphere. Spot groups in one hemisphere, i.e., either north or south of the Sun's equator, tend to be oriented similarly, with the same polarity on the leading edge. Spot groups in the other hemisphere are oriented predominantly in the opposite direction. Spots last for a few days to weeks, with subsequent spots often appearing at similar longitudes.

The number of spots visible on the Sun varies over a roughly 11-year cycle. Early in the cycle there are relatively fewer spots and they erupt at intermediate latitudes, roughly 30 degrees or so. As the cycle progresses, the number of spots on the Sun increases and the latitudes of those spots gradually migrate toward the equator. The Sun may go from having no spots for weeks on end to having 200 or so near the peak of the activity cycle. Still later in the cycle the spot numbers decrease again and spots get even closer to the equator. Spot numbers often rise faster than they fall, but every cycle is a bit different both in terms of duration and of the number of spots present at maximum. High-latitude coronal activity and the first few spots indicative of the beginning of the next activity cycle are often apparent at least two years before the end of a current cycle. When the next cycle's spots appear we see that the Sun's magnetic pole will have flipped and the N-S magnetic orientation of the spot groups will have reversed hemispheres.

The following two plots show the areal coverage and latitude distribution of sunspots since the mid-1870s. The top plot is called a butterfly diagram, for obvious reasons.

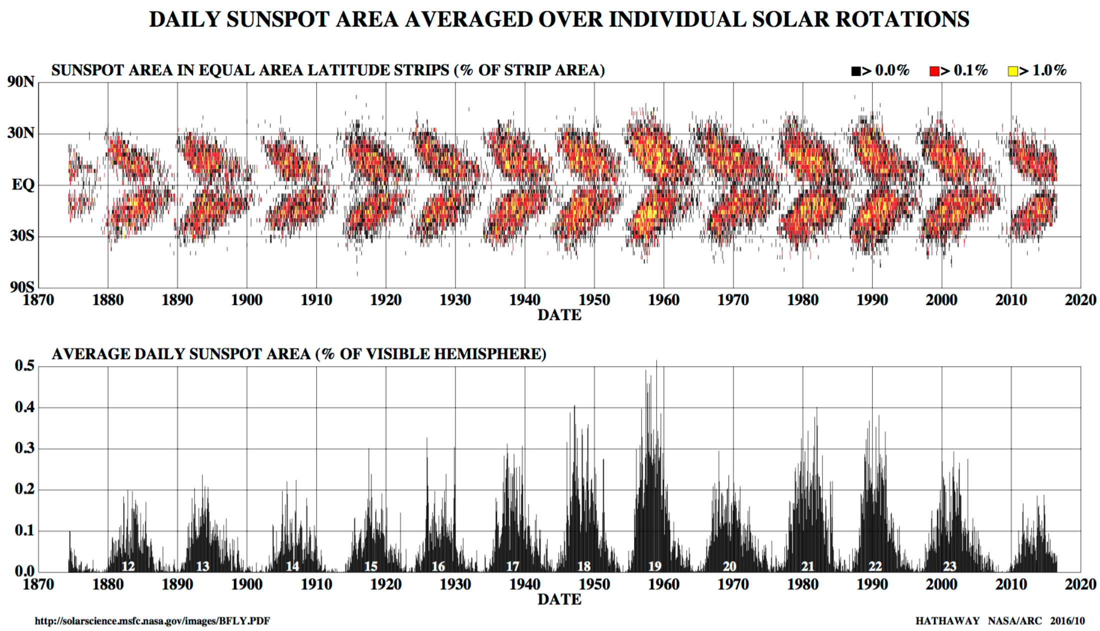


Figure 13.21: a) butterfly diagram and b) daily sunspot area. http://solarscience.msfc.nasa.gov/SunspotCycle.shtml

By the mid-1800s enough observations of sunspots existed that this cyclic behavior was clear. Reconstructions of sunspot numbers over the previous century showed a definite rise and fall between 1755-1766, and that cycle is cycle number 1. The plot above picks up at the tail end of cycle 11. You can see from the lower plot in the figure above that the amount of the solar disk that is covered by spots varies considerably from one solar maximum to the next. In fact, there have been times when a “cycle” had no spots at all. The first such extended minimum known, from ~1645 to ~1715, is called the Maunder Minimum, which was identified in the late 1800s by Annie Maunder and husband Walter.

It is possible to push evidence for the level of solar magnetic activity back for approximately 11,000 years. One of the principal techniques is to measure the level of ^{14}C in tree rings. Carbon-14 is radioactive, with a half life of 5,730 years. It forms in the atmosphere when neutrons created by energetic cosmic rays interact with ^{14}N. When the Sun is more active the solar wind is stronger; when the solar wind is stronger, more galactic cosmic rays are deflected and prevented from entering the inner solar system. When the Sun is less active, more ^{14}C is produced and more $^{14}CO_2$ is incorporated into trees. The mid-late twentieth-century Sun was more active than its average level over the last 11,400 years.

When the Sun is at the peak of the magnetic activity cycle there are more spots on the surface and you might expect that the Sun would be overall less luminous. Although it is a bit counterintuitive, the Sun is actually more luminous at periods of high activity. The spots are dark, yes, and a large spot group crossing the disk of the Sun will make a noticeable dip in the solar irradiance but that dip due to the spots occurs from a higher-than-average background. On average, we receive ~1,366 $J / (s \cdot m^2)$ from the Sun; at solar maximum the Sun is ~ 1 $J / (s \cdot m^2)$ brighter than at minimum, or ~0.07%. Yes, the Sun is a variable star.

Although a more active than average Sun is brighter, it isn’t responsible for the current warming climate trends. That said, the Sun’s activity can have some effects. One classic example, though, has lately been disputed: during the Maunder Minimum temperatures in Western Europe were slightly but noticeably colder than normal, leading that time period to be called a “little ice age”, and climate scientists have argued whether or not the less active Sun played a role or whether the cooling might more reasonably be attributed to volcanic activity. Whether the solar irradiance variation plays a role in the weather patterns such as La Niña and El Niño, that result from temperature variations in the surface waters of the eastern Pacific Ocean, is a subject of active research.

Magnetic activity on other stars

The Sun's differential rotation and outer convective zone combine to sustain a magnetic dynamo that models suggest operates at the tachocline at the base of the convective zone. In the Sun's atmosphere, magnetic activity is manifest as sunspots at the photospheric level through prominences, flares, coronal mass ejections, etc., farther out in the atmosphere. The existence of the chromosphere and the hot corona, with its varying structure including streamers and coronal holes, seems to be powered primarily by the magnetic activity. Do other stars show similar or related magnetic activity? For many the answer is yes.

As we will see in the chapters on stellar spectra and evolution, stars that, like the Sun, have modest photospheric temperatures also have outer convective zones. If we look simply at stars that are producing their energy by hydrogen fusion, thin outer convective layers appear in stars of roughly 1.5 solar masses, with surface temperatures of roughly 7,000 K; convective zones get steadily deeper as we move down through stars of lower temperatures, and stars smaller than about 0.3 solar masses, or having surface temperatures less than about 3,000 K, are fully convective. Higher-mass stars will have outer radiative zones for the majority of their lives but their surfaces will expand and cool as the stars evolve to become red giants or supergiants, and at that point they, too, will develop outer convective zones. The nature of the magnetic dynamo may change somewhat as stars become fully convective, but the change must be smooth because stars across the full range of outer convective zone depth tend to show some level of magnetic activity.

Stars form with a range of angular momenta but generally tend to rotate more rapidly when they are young. For instance, a solar-mass star in the Pleiades, aged roughly 125 Myr, might rotate in a few days rather than a few weeks. Two main processes act to slow rotation. Solar-type stars with magnetic activity will have a stellar wind and, like the Sun, experience angular momentum loss as charged particles in the wind are forced by the magnetic field to co-rotate with the stellar surface out to large distances in the process of *magnetic braking*. Regardless of its initial mass, as a star evolves to become a giant its core will shrink somewhat and its outer envelope expand tremendously, with the net result being that the star's moment of inertial will increase, also leading to slower rotation. Slower rotation results in less rotational energy to power a magnetic dynamo. Among stars with outer convective zones and evidence of magnetic activity, younger, more rapidly rotating stars tend to be more active, although for the fastest rotators activity tends to saturate.

How do we know that stars are active? Measuring other stars' magnetic fields directly is not trivial. Spatially resolved images of stars' surfaces are not yet routinely possible, although for some stars — usually giants, usually fairly close — large telescopes have been able to detect surface brightness variations and it may eventually be possible to categorize a range of types of granulation patterns or starspot patterns. Even without the spatial resolution, though, like the Sun, magnetically active stars may show evidence of flaring, or have x-ray emission suggestive of the existence of a corona. One prominent signature of magnetic activity, measurable from ground-based observatories, is the strength of the emission cores in the H and K spectral absorption lines of singly ionized calcium (Ca II H+K). Let's unpack that a bit. The interior of a star produces a roughly blackbody spectrum. The photosphere of the star, being cooler, introduces absorption lines into the spectrum: photospheric atoms, with their electrons in relatively lower energy levels, absorb particular wavelengths of light that correspond to their permitted energy transitions. On the other hand, a relatively hot gas, whether or not it's in front of a blackbody, will produce emission lines as excited electrons drop to lower energy levels. A star like the Sun has a photosphere and nearby temperature minimum, a chromosphere and transition region in which the temperature rises again, and a hot corona. Some spectral lines are formed by atoms across a range in height over which those spectral lines change from absorption to emission (and possibly back again; which process wins, absorption or emission, isn't always simple or straightforward). Ca^+ has two such lines, called H and K for historical reasons; the H and K lines have central wavelengths of 396.8 and 393.4 nm, respectively. These lines are very prominent, broad, absorption lines in the spectra of cool stars and even the most inactive stars show some basal level of core line emission. In the Sun, the emission in the line cores is stronger over regions where the magnetic field is stronger. The strength of the average line-core emission component is correlated with the star's luminosity — this is called the Wilson-Bappu Effect — so astronomers studying the level and variation in stellar magnetic activity tend to measure the ratio of the emission feature strength to the average level of the nearby continuum (i.e., outside the line) spectrum. The active regions for unsaturated stellar surfaces are carried across the surface as the star rotates, so an added benefit of studying the Ca II H+K emission (or other similar indicators) is acquiring an independent measurement of a star's rotation rate. The figure below shows roughly what the center of the Ca II K line looks like for an active solar-like star.

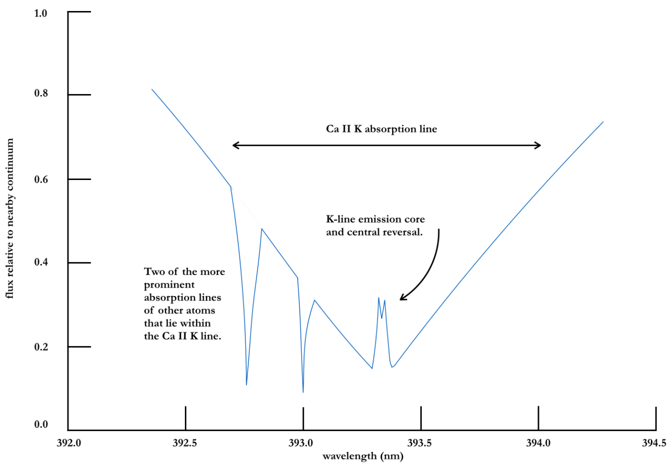


Figure 13.22: Sketch of the emission core of the Ca II K line for a magnetically active star.

The Ca II H+K lines are not the only absorption lines to show central emission feature that is sensitive to the level of stellar magnetic activity, but, being both strong and in the visible region of the spectrum, they are among the most straightforward to study. In the 1960s astronomer Olin Wilson (of the Wilson-Bappu Effect) began a systematic program at the Mount Wilson Observatory of observing Ca II H+K emission levels for cool stars; the program he launched ultimately produced data spanning three decades for some of the target stars. At their most active, young stars' surfaces seem to be saturated with dark spots and these stars will be a bit less luminous when their Ca II H+K emission is highest. Older stars are more like the Sun, i.e., a bit brighter when most active despite having some regions with cooler, darker spots. Young stars (i.e., younger than about 200 million years) also tend to be a bit more erratic, varying strongly but not necessarily cyclically; many older stars show evidence of some multi-year periodicity in the variation of their Ca II H+K emission that may be akin to our sunspot cycle. Even among moderately slowly rotating stars like the Sun there seems to be considerable variation in the levels of activity, with some stars being several times more variation due to starspots than the Sun currently does. Some, particularly older, stars are stable at least at the level of ~0.5%. Eventually the oldest, slowest rotators should reach a point at which they no longer have sufficient rotational energy to power a magnetic dynamo, although recalling that the Sun has experienced extended periods of low activity, such as the Maunder Minimum, we can't conclude that any given star has no magnetic activity.

Sample problems

1. The Sun's effective temperature is 5,778 K (a star's T_{eff} is the temperature of a blackbody with the same total flux as our star) and its radius is ~695,000 km. Calculate the "solar constant", i.e., the amount of energy received per second per square meter by a detector at 1 A.U. (but above our atmosphere) oriented perpendicular to the incoming sunlight. The Sun's luminosity isn't quite contant, but you may assume that it is for purposes of this problem.

2. The principal fusion process at work in the Sun is the proton-proton chain. Explain what role each of the four known fundamental forces plays in the proton-proton chain. Recall that there's more info about forces and fusion reactions in the introductory chapter.

3. The Sun will use about 10% of its mass during the hydrogen-fusing phase of its life. In Ch. 1, problem 11, we saw that each reaction in which 4 hydrogens fuse into 1 helium releases $4.3 \cdot 10^{-12}$ J. Given the Sun's mass and luminosity, calculate approximately how long the Sun could do hydrogen fusion. Assume the over this period the Sun's luminosity is constant. . .which isn't true, but for purposes of this problem it isn't too wildly wrong.

4. Sketch a cross-sectional view of the Sun and label the various regions.

5. Reading carefully? Briefly explain

a) solar wind
b) coronal hole
c) solar neutrino problem
d) helioseismology
e) granulation
f) limb darkening
g) filament vs. prominence
h) coronal mass ejection
i) solar flare
j) active region
k) Maunder minimum
l) differential rotation
m) plasma
n) coronagraph
o) magnetic braking
p) forbidden transition

5. Check out a few web sites for information about the current conditions of Sun:

a) At http://www.spaceweather.com find the speed and density of the solar wind at the Earth.
b) Same spaceweather web site: how many days, if any, has it been since there were last sunspots visible?
c) At https://sdo.gsfc.nasa.gov/data/ you should see several images of the Sun; is there an obvious coronal hole? (the 193 and 211 Å) images, near the top of the page are particular good for this.
d) At this SDO web site scroll down the page until you find the Dopplergram; what is this image showing?
e) At https://www.swpc.noaa.gov you should see images and graphs of the solar x-ray flux, proton flux, and prediction for the current auroral oval showing where aurorae might be visible. Have there been any CMEs in the last few hours?

Answers to selected problems are on the next page:

1. One way to tackle this: Use $4\pi r^2 \sigma T^4$ to calculate the luminosity of the Sun (and check your value with that given at the beginning of the chapter to make sure you did it correctly). Divide that value by the surface area of a sphere of radius 1 AU, being careful to convert 1 AU into meters. A shortcut would be to note that there's a factor of 4π in both the numerator and denominator of this ratio and cancel both of those factors of 4π before doing the calculation. You should get about 1,360 J/m^2 for your answer.

3. One way to tackle this: follow the units. [(J/s) / (J/rxn)] x (kg of H used / rxn) gives you kg / sec hydrogen used to maintain the Sun's luminosity. Total kg available / (kg / sec) gives you total seconds. When you convert that to years, you should get about 10^{10} years.