The Solar Atmosphere

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The sunlight we usually see is the light in the visible spectrum (white light). Here we are seeing the photosphere.
Eclipse Images

But if we can block out the light from the photosphere, different parts of the atmosphere become visible.
The Layers of the Solar Atmosphere

[Diagram showing the temperature and density distribution across the solar atmosphere, with labels for the photosphere, chromosphere, and corona.]

- Solar photosphere
- Solar chromosphere (courtesy of J. Okamoto NAOJ)
- Solar corona (AIA SDO)
The light that comes from the Sun

The both the intensity and wavelength distribution of light we receive from the Sun follows to a good degree the blackbody emission at 6000K

From Green (1982)
The Visible Solar Spectrum

Of the visible light emitted by the Sun, not all of it can escape, some is absorbed by atoms (and sometimes molecules) in the atmosphere (see dark lines in the spectrum below).

Above is the visible light spectrum. The dark lines show the absorption of light by elements in the solar atmosphere.

Helium was discovered from spectral observations of the Sun by Jules Janssen in 1868 and Norman Lockyer (who realized it was a new element).
The radiation in the corona is generally optically thin (meaning you can ignore effects that come from the optical depth).

Photospheric and chromospheric plasma is dense enough to reabsorb some of the radiation.
Spectral lines as a way to investigate different heights in the atmosphere

From Vernazza, Avrett and Loesner (1981)
Some other properties that can be important

VAL-C model + smooth extrapolation to the corona

- e-n collision dominates in lower chromosphere (⇒ weak ionization)
- e-i collision dominates in upper chromosphere (⇒ partial ionization)
The elemental abundances in the corona differ from those of the photosphere, but only for elements with lower FIP.

**Table I. First I.P. and photospheric abundances of the abundant solar elements**

<table>
<thead>
<tr>
<th>Element</th>
<th>I.P. (eV)</th>
<th>log$_{10}$ Abundances</th>
<th>Abundances relative to O</th>
<th>Abundances relative to Mg</th>
</tr>
</thead>
<tbody>
<tr>
<td>H</td>
<td>13.6</td>
<td>12.00</td>
<td>1170</td>
<td>26300</td>
</tr>
<tr>
<td>He</td>
<td>24.6</td>
<td>10.99</td>
<td>115</td>
<td>2570</td>
</tr>
<tr>
<td>C</td>
<td>11.3</td>
<td>8.60$^4$</td>
<td>0.47</td>
<td>11.0</td>
</tr>
<tr>
<td>N</td>
<td>14.5</td>
<td>8.00$^3$</td>
<td>0.12</td>
<td>2.6</td>
</tr>
<tr>
<td>O</td>
<td>13.6</td>
<td>8.93</td>
<td>1.0</td>
<td>22.0</td>
</tr>
<tr>
<td>Ne</td>
<td>21.6</td>
<td>8.11$^2$</td>
<td>0.15</td>
<td>3.4</td>
</tr>
<tr>
<td>Na</td>
<td>5.1</td>
<td>6.33</td>
<td>0.0025</td>
<td>0.056</td>
</tr>
<tr>
<td>Mg</td>
<td>7.6</td>
<td>7.58</td>
<td>0.045</td>
<td>1.0</td>
</tr>
<tr>
<td>Al</td>
<td>6.0</td>
<td>6.47</td>
<td>0.0035</td>
<td>0.078</td>
</tr>
<tr>
<td>Si</td>
<td>8.2</td>
<td>7.55</td>
<td>0.042</td>
<td>0.94</td>
</tr>
<tr>
<td>S</td>
<td>10.4</td>
<td>7.21</td>
<td>0.019</td>
<td>0.43</td>
</tr>
<tr>
<td>Ar</td>
<td>15.8</td>
<td>6.65$^3$</td>
<td>0.0053</td>
<td>0.12</td>
</tr>
<tr>
<td>Ca</td>
<td>6.1</td>
<td>6.36</td>
<td>0.0027</td>
<td>0.060</td>
</tr>
<tr>
<td>Fe</td>
<td>7.9</td>
<td>7.51$^3$</td>
<td>0.038</td>
<td>0.85</td>
</tr>
<tr>
<td>Ni</td>
<td>7.6</td>
<td>6.25</td>
<td>0.0021</td>
<td>0.047</td>
</tr>
</tbody>
</table>

1 Adopted from Anders and Grevesse [2]. Abundances are given in units in which the log$_{10}$ abundance of H is fixed at 12.

**Fig. 2.** A schematic representation of the SW and SEP abundances relative to those in the photosphere. Values are plotted against the First Ionization Potential (FIP)
Magnetic field strength throughout the atmosphere

From Gary (2001)
How to determine other important properties of the atmosphere

MHD momentum Equation

\[ \rho \frac{\partial \mathbf{v}}{\partial t} + \rho \mathbf{v} \cdot \nabla \mathbf{v} = -\nabla p + \mathbf{J} \times \mathbf{B} + \rho \mathbf{g} \]

The size of each term

\[ \frac{\rho v^2}{L} = \frac{P}{L} + \frac{B^2}{\mu_0 L} + \rho g \]

We can work out which terms are important for modelling different layers of the solar atmosphere by comparing the size of the different terms and neglecting small terms.
Pressure scale height: $\Lambda$

Comparing terms 2 and 4

$$\frac{P}{L}, \rho g$$

The pressure scale height is the height over which the pressure in the atmosphere drops by a factor of $e^1$.

Using hydrostatic equilibrium

$$\frac{\partial P}{\partial z} = -\rho g$$

And an ideal gas

$$P = \frac{\rho RT}{\mu}$$

Then we have

$$\frac{\partial P}{\partial z} = -P \frac{\mu g}{RT} = -P \frac{1}{\Lambda}$$

If $L \ll \frac{P}{\rho g} = \Lambda$ then it may be acceptable to neglect the gravity term.
Plasma Beta $\beta$

(As seen in the MHD lecture) Comparing terms 2 and 3

\[ \frac{P}{L}, \frac{B^2}{\mu_0 L} \]

The plasma beta is the comparison between the gas pressure and magnetic pressure.

The magnetic pressure is given by: \( \frac{B^2}{2\mu_0} \) and the gas pressure by: \( P \)

\[ \beta = \frac{P}{B^2/2\mu_0} \]

If \( \beta \gg 1 \) then

\[ \rho \frac{\partial \mathbf{v}}{\partial t} + \rho \mathbf{v} \cdot \nabla \mathbf{v} = -\nabla p + \rho \mathbf{g} \]

If \( \beta \ll 1 \) then

\[ \rho \frac{\partial \mathbf{v}}{\partial t} + \rho \mathbf{v} \cdot \nabla \mathbf{v} = \mathbf{J} \times \mathbf{B} + \rho \mathbf{g} \]
Plasma Beta $\beta$ throughout the atmosphere

Switch from high to low beta happens in the chromosphere

From Gary (2001)
Mach numbers

This gives information on how fast the flow is moving compared to the speed at which information is being transmitted.

Comparison between terms 1 and 2 (hydrodynamic Mach number)

\[ M^2 = \frac{v^2}{c_s^2} = v^2 \left( \frac{\gamma P}{\rho} \right)^{-1} \]

(note I have used the square as this is a good measure of compressibility)

Comparison between terms 1 and 3 (Alfvénic Mach number)

\[ M_A^2 = \frac{v^2}{V_A^2} = v^2 \left( \frac{B^2}{\mu_0 \rho} \right)^{-1} \]

If these are small then \( \nabla \cdot \mathbf{v} = 0, \frac{\partial \mathbf{v}}{\partial t} = 0 \) or \( \mathbf{v} = 0 \) may be appropriate simplifications
The photosphere

This is the base of the solar atmosphere and is defined as the height at which the visible light photons are likely to escape. The majority of the light that escapes from the Sun come from photons that are emitted from the photosphere.
Vertical photospheric magnetic field

Hinode SOT magnetogram saturated at above 500G (0.05T). The strong magnetic field has collected at the boundaries of supergranules.
Hinode SOT magnetogram saturated 100G (0.01T). The magnetic field is beginning to fill the regions between the supergranule boundaries.
Vertical photospheric magnetic field

Hinode SOT magnetogram saturated 25G (0.0025T). The magnetic field is now inside granules.
Photospheric magnetic field

Horizontal magnetic field observed by Hinode SOT magnetogram saturated 200G (0.02T). Observed near the edge of granules.
Why do we have these field strengths?

1) We can use arguments based on energy equipartition

\[
\frac{1}{2} \rho v^2 \sim \frac{B^2}{2\mu_0} \rightarrow \quad B \sim \sqrt{\mu_0 \rho v} = \sqrt{4\pi \times 10^{-7} \times 2 \times 10^{-4} \times 2 \times 10^3} = 0.02 \text{T (200 G)}
\]

2) By equating the gas pressure of the photosphere \( (p_e) \) to the magnetic pressure inside the magnetic element:

\[
p_e \sim \frac{B^2}{2\mu_0} \rightarrow \quad B \sim \sqrt{2\mu_0 p_e} = \sqrt{8\pi \times 10^{-7} \times 5 \times 10^3} \sim \sqrt{10^{-2}} = 0.1 \text{T (10}\,^3\, \text{G)}
\]
Sunspots

Drawing by John of Worcester on 8/12/1128 showing two spots on the Sun. This is believed to be the oldest drawing of a Sunspot in existence.

https://sunearthday.nasa.gov/2006/locations/firstdrawing.php
Sunspot observations before modern observations

Sunspot sketches by Galileo Galilei from 1612 (taken from rice.edu)

Sunspot number since the 1600s
Emerging magnetic flux to give sunspot formation

The complex flow inside the Sun stretch and fold the magnetic field until it becomes strong enough to rise up through the solar surface and form Sunspots.
An image of a sunspot (left) with its magnetic field measurement (right). Field strengths in the sunspot are typically $10^3$G ($0.1$T is SI units).
The photosphere: granulation

Granules are cells of convection that can be seen on the surface of the Sun. Characteristic flow speeds of a few km/s.

Power spectra

Matsumoto & Kitai (2010)
Because the plasma $\beta$ has become sufficiently small, the magnetic field structures the chromosphere.
How we define the Chromosphere

Chromosphere, meaning sphere of colour, describes the layer of the solar atmosphere that is clearly visible in a variety of spectral lines in the visible, UV and IR wavelengths.
The spectral lines of the Chromosphere

These are three very important spectral lines for investigating the chromosphere.

From Vernazza, Avrett and Loesner (1981)
The spectral lines of the Chromosphere

From De Pontieu et al (2014)
Chromospheric dynamics: Spicules

Spicules are thin jets of chromospheric plasma that shoot up through heights of $O(5 \times 10^6 \text{m})$. They have speeds of $O(50 \text{ km/s})$ and generally follow ballistic motion.

Recently a second type of spicule (type II spicules) was discovered that have a different thermodynamic (and possibly dynamic) evolution.
How are neutrals coupled to magnetic fields

① Movement of magnetic field
② together with charged particles
③ Charged particles collide with neutrals
④ Neutrals move in the same direction as the charged particles
To understand the equations used in partially ionised MHD it helps if they are derived. Straying with a neutral and an ionised fluid

Neutral Fluid:

\[ \rho_N \frac{Dv^N}{Dt} = -\nabla p_N + \rho_N \mathbf{g} - C_I \]

Ionised Fluid:

\[ \rho_I \frac{Dv^I}{Dt} = -\nabla p_I + \mathbf{J} \times \mathbf{B} + \rho_I \mathbf{g} - C_N \]

Where represents C the terms that arise from collisions between ions and neutral give as (note they are equal in magnitude but opposite in sign)

\[ C_N = \rho_I v_{IN} (v_I - v_N) \quad \quad C_I = \rho_N v_{NI} (v_N - v_I) \]

\[ \rho_I v_{IN} = \rho_N v_{NI} = \rho_N \rho_I \frac{1}{2m_i} \sqrt{\frac{8k_B T}{\pi m_i}} \Sigma_{in} \quad \Sigma_{in} \sim 5 \times 10^{-15} \text{ cm}^2 \]
Partially Ionised MHD

Assuming that $\rho_i \ll \rho_n$ and $p_i \ll p_n$ then comparing the two underlined terms:

$$\rho_I \frac{Dv_I}{Dt} = -\nabla p_I + \mathbf{J} \times \mathbf{B} + \rho_I \mathbf{g} + \rho_N \nu_{NI} (\mathbf{v}_N - \mathbf{v}_I)$$

Then we can take that:

$$\rho_I \frac{Dv_I}{Dt} = -\nabla p_I + \mathbf{J} \times \mathbf{B} + \rho_I \mathbf{g} + \rho_N \nu_{NI} (\mathbf{v}_N - \mathbf{v}_I)$$

This is known as strong coupling
Partially Ionised MHD

Now comparing the three force terms:

\[ \rho_1 \frac{Dv_I}{Dt} = -\nabla p_I + \mathbf{J} \times \mathbf{B} + \rho_I \mathbf{g} + \rho_N \nu_{NI} (\mathbf{v}_N - \mathbf{v}_I) \]

If \( \beta \) is relatively small as \( \rho_i \ll \rho_n \) and \( p_i \ll p_n \) then :

\[ \rho_1 \frac{Dv_I}{Dt} = -\nabla p_I + \mathbf{J} \times \mathbf{B} + \rho_I \mathbf{g} + \rho_N \nu_{NI} (\mathbf{v}_N - \mathbf{v}_I) \]
Partially Ionised MHD

Now we have:

\[ \mathbf{J} \times \mathbf{B} = -\rho_N \nu_{NI} (\mathbf{v}_N - \mathbf{v}_I) \]

\[ \rho_N \frac{Dv_N}{Dt} = -\nabla p_N + \rho_N \mathbf{g} + \mathbf{J} \times \mathbf{B} \]

i.e. the ideal MHD momentum equation, but for the neutral fluid velocity.
Partially Ionised MHD

The induction equation now has to be revised to remove the ion velocity:

$$\frac{\partial \mathbf{B}}{\partial t} = \nabla \times \left[ \mathbf{v}_I \times \mathbf{B} - \eta \mathbf{J} \right]$$

Using the ionised fluid equation of motion, it can be derived that:

$$\mathbf{v}_I = \mathbf{v}_N + \frac{\mathbf{J} \times \mathbf{B}}{\rho_N \nu_{NI}}$$

So, the induction equation becomes:

$$\frac{\partial \mathbf{B}}{\partial t} = \nabla \times \left[ \mathbf{v}_N \times \mathbf{B} + \frac{(\mathbf{J} \times \mathbf{B}) \times \mathbf{B}}{c \rho_N \nu_{NI}} - \eta \mathbf{J} \right]$$

Where $$\frac{(\mathbf{J} \times \mathbf{B}) \times \mathbf{B}}{\rho_N \nu_{NI}}$$ is the ambipolar diffusion term.
The Corona
Viewing the corona

Visible (white) light

Radio

EUV

X-ray

2000/06/05
How does the light we can’t see vary over a solar cycle

Low solar activity

High solar activity

It is clear that the total amount of X-ray photons at high solar activity is much greater than at low activity
What about the Coronal magnetic field

The topology and strength of the magnetic field of the Sun, and with it the heliosphere, changes drastically over a cycle. Active regions create more regions of closed field, and weaken the dipole moment.

From nasa.gov
Coronal structure

The corona is made up of:
- Coronal holes (open magnetic field regions)
- Quiet regions (bright points, microflares, nanoflares)
- Loops
- Active regions
How do you get the 1MK Corona?

The solar corona seen in x-ray observed by Yohkoh
What heating mechanisms can exist?

The energy source for the heating is the magnetic field, but the dissipation is a story of waves VS reconnection (or AC vs DC).

Energy associated with motion in the lower atmosphere can be transported by waves into the corona and dissipated there.

Current sheets of a tangled field reconnect, releasing energy and heat the corona.

Parker (1983)
Coronal loops are one of the fundamental structures in the corona tracing the magnetic field in the corona. Length $O(5 \times 10^7 \text{m})$ and width $O(10^6 \text{m})$ (though many different sizes exist!)
Thermal conduction

To have the clear structuring along the magnetic field that is observed for coronal loops, we need a process that makes the temperature along the loop relatively uniform. This process is thermal conduction.

Energy equation for the corona

\[
\frac{\rho^\gamma}{\gamma - 1} \frac{D}{Dt} \left( \frac{\rho}{\rho^\gamma} \right) = \nabla \cdot (\kappa \nabla T) - \rho^2 Q(T) + \frac{j^2}{\sigma} + H
\]

- Thermal conduction
- Joule heating
- Radiative losses
- Other heating terms

Thermal conduction is about $10^{13}$ times stronger along the direction of the field in the corona than across it.
Solar flares

Flares are characterized by the impulsive intensity increase in multiple wavelengths. An example of the soft x-ray (related to the thermal emission) is shown below.

A movie of a flare from EUV observations using the TRACE satellite. The bright structures you can see at the end are called post-flare loops.

Timescale: few hours
Temperature: $\sim 3 \times 10^7$ K
Energy: $10^{16}$ to $10^{18}$ Joule
Non-thermal flare emission

The CSHKP Model

The key process to explain flares is magnetic reconnection.

X8-class flare on September 10, 2017 SDO/AIA 211

Carmichael 1964; Sturrock 1966; Hirayama 1974; Kopp & Pneuman 1976

Post-flare loop

FMSS

plasmoid

reconnection

post-flare loop

inflow

outflow jet

magnetic polarity inversion line (PIL)

flare ribbon

McKenzie 2002

SDO/AIA 211

16:09:49 UT Sep 10, 2017

SDO/AIA 211

15:39:35 UT Sep 10, 2017

15:54:35 UT Sep 10, 2017
Active Region 10930

Observation of flare ribbons in the lower atmosphere

Before the flare, the sunspots rotate, building up magnetic energy
With Complex Magnetic Field

Not only can the rotation build up magnetic energy, but the complex magnetic structure tells us that there is lots of current in the system, making it ripe for magnetic reconnection.

This small-scale magnetic structuring can provide a trigger for flares to start (Kusano et al. 2012).
Prominences and Filaments

11 August 1980: Hα image

Source: NOAA/SEL/USAF
Prominences across different wavelengths

From Okamoto et al (2015)
Quiescent filament observed in Hα line centre on 21st Dec 2010 at Hida observatory

Height: 10⁷ to 10⁸ m
Length: 6x10⁷ to 6x10⁸ m
Width: 5x10⁶ to 2x10⁷ m
(Tandberg-Hanssen 1995)
Prominences and the photospheric magnetic field

Prominences always lie above the polarity inversion line.
Prominence magnetic field

Field strength of 15G
Field inclination $\theta_B$
Field azimuth $\chi_B$

The field strength is about 10-15 Gauss and increases up to 40 G: projection effects (?)
The field is 90 degree inclined with respect to the local vertical in the prominence body and becomes slightly more vertical at the “barb”

Prominence magnetic field is horizontal

Horizontal field
30° to line of sight
Prominences & filaments: Basic model

You’ve already been introduced to the Kippenhan-Schluter model for prominences (c.f. MHD lecture by Browning on Monday)

Prominence mass collects in dips in the magnetic field. And magnetic tension supports the mass against gravity.
Quiescent prominence

With the launch of Hinode SOT, we now know that quiescent prominences can be highly dynamic, full of nonlinear flows, waves and instabilities, basically a turbulent medium.

Dominated by field aligned flows, rotations (e.g. Okamoto et al 2016) and MHD waves (e.g. Okamoto et al 2007). Clearly magnetically dominated.
Why the difference?

The size of the tension term
\[ \frac{B_h \partial B_z}{\mu_0 \partial h} \sim \frac{1}{2\mu_0} \frac{\partial B^2}{\partial h} \sim \frac{B^2}{2\mu_0 L} \]

The size of the gravity term
\[ \rho g = \frac{P}{\Lambda} \]

Comparing the size of these terms we get (I use TG here for want of a better name):
\[ TG = \frac{P}{B^2 / 2\mu_0} \frac{L}{\Lambda} = \beta \frac{L}{\Lambda} \]

<table>
<thead>
<tr>
<th></th>
<th>$\beta$</th>
<th>$L(m)$</th>
<th>$\Lambda(m)$</th>
<th>$TG$</th>
</tr>
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<tbody>
<tr>
<td>Quiescent</td>
<td>0.1</td>
<td>$5 \times 10^7$</td>
<td>$4 \times 10^5$</td>
<td>$\sim 10$</td>
</tr>
<tr>
<td>Active region</td>
<td>0.005</td>
<td>$10^7$</td>
<td>$4 \times 10^5$</td>
<td>$\sim 0.1$</td>
</tr>
</tbody>
</table>

Note that this is somewhat a crude metric
How do we have cool material in the hot corona?

A plot of the radiative losses in the solar corona for a fixed pressure across a wide range of temperatures.

If the plasma can start to cool, it can become thermally unstable. When this happens the next stable temperature is near $10^4$K where optically thick radiation becomes important.

From Anzer and Heinzel (2008)
Prominences & filaments: Eruptions

What triggers the eruption, be it ideal MHD instability, flux emergence or reconnection, is still an active research topic.
### Basic physical quantities

<table>
<thead>
<tr>
<th></th>
<th>Photosphere</th>
<th>Chromosphere</th>
<th>Corona</th>
</tr>
</thead>
<tbody>
<tr>
<td>( \rho )</td>
<td>( 2 \times 10^{-4} )</td>
<td>( 10^{-6} ) to ( 10^{-9} )</td>
<td>( 2 \times 10^{-12} )</td>
</tr>
<tr>
<td>( T ) (K)</td>
<td>( 4.5 \times 10^3 ) to ( 6 \times 10^3 )</td>
<td>( 5 \times 10^3 ) to ( 2 \times 10^4 )</td>
<td>( \geq 10^6 )</td>
</tr>
<tr>
<td>( B ) (T)</td>
<td>0.001 to 0.3</td>
<td>0.001 to 0.1</td>
<td>0.0001 to 0.003</td>
</tr>
<tr>
<td>( L ) (m)</td>
<td>( 4 \times 10^6 ) to ( 2 \times 10^7 )</td>
<td>( 4 \times 10^6 ) to ( 2 \times 10^7 )</td>
<td>( 10^6 ) to ( 5 \times 10^7 )</td>
</tr>
<tr>
<td>( V ) (km/s)</td>
<td>0.1 to 3</td>
<td>2 to 100</td>
<td>10 to 100</td>
</tr>
<tr>
<td>( \Lambda ) (m)</td>
<td>( 2 \times 10^5 )</td>
<td>( 4 \times 10^5 )</td>
<td>( 5 \times 10^7 )</td>
</tr>
<tr>
<td>( \beta )</td>
<td>1 to 1000</td>
<td>10 to 0.01</td>
<td>0.1 to 0.0001</td>
</tr>
<tr>
<td>( v_A ) (km/s)</td>
<td>0.05 to 7</td>
<td>30 to 50</td>
<td>&gt;300</td>
</tr>
<tr>
<td>( c_s ) (km/s)</td>
<td>7</td>
<td>10</td>
<td>100</td>
</tr>
<tr>
<td>( \xi_i )</td>
<td>( 10^{-3} ) to ( 10^{-6} )</td>
<td>( 10^{-1} ) to ( 10^{-6} )</td>
<td>1</td>
</tr>
<tr>
<td>Electron Mean free path (m)</td>
<td>( 10^{-5} )</td>
<td>( 10^0 )</td>
<td>( 10^7 )</td>
</tr>
</tbody>
</table>

Note that these often depend on proximity to an active region
The messy solar atmosphere
Conclusion

• Solar Atmosphere:
  • Photosphere to Chromosphere through the Transition region to the Corona

• Host to a vast array of plasma regimes:
  • Temp increases by a factor of 200 or more
  • Density decreases by about $10^8$

• Processes occur on a wide array of length and time scales
  • Lengths: electron mean free path ($10^{-5}\text{ m}$) up to global solar radii (100 Mm)
  • Times: fraction of a second up to days/months