Planet formation

Astro 9601
Topics to be covered

- Observational constraints (13.1)
- Star formation
  - Observations of star formation (13.3.3)
- Evolution of the solar nebula (13.4)
- Condensation of solid bodies (13.5)
  - note: I take a somewhat different tack here from DeP&L
- Formation of the terrestrial planets (13.6)
- Formation of the giant planets (13.7)
- Planetary migration (13.8)
- Small bodies in orbit around the Sun (13.9)
- Origin of planetary satellite systems (13.11)
Formation of the Elements

• Cosmological Nucleosynthesis:
  • Most H and He formed shortly (few minutes) after the Big Bang
  • H and He make up most of the universe

• Stellar Nucleosynthesis:
  • Most elements heavier than H and He (astronomical “metals”) form either in red giants or supernovae by nuclear fusion.
  • Those metals released into the interstellar medium by supernovae, then mix with the H/He clouds, and get incorporated into the next round of star formation if/when these clouds collapse.

SN remnants (HST)
Star-forming region: Orion Nebula

- Distance = 1,500 ly
- Age < 1 million yrs
- Young stars:
- Nearest massive star-forming region

near-IR
(more sources visible than in optical image below)
Orion Nebula: protoplanetary disks seen in silhouette only

250-1000 AU diameters
Solar System Formation: overview

Cloud of interstellar gas (the **solar nebula**) begins to contract, condense, and spin faster through conservation of angular momentum.

~10,000 AU across

~1,000 AU across
Gas and dust particles collide and coalesce in the proto-planetary disk (~100 AU across)

Multiple collisions and accretion into planets (~40 AU across)
Our Solar System
Terrestrial Planets: Mercury, Venus, Earth, and Mars

- Rocky (silicate) outer parts (crust and mantle) and inner cores of Fe-Ni metal.
- Silicates are compounds of silica ($\text{SiO}_2$) and the oxides of other metals.
- Common silicates:
  - Olivine: $(\text{Mg,Fe})\text{SiO}_4$
  - Pyroxene: $(\text{Mg,Ca,Fe})_2\text{Si}_2\text{O}_6$
  - Feldspar (e.g., $(\text{Na,K})\text{AlSi}_3\text{O}_8$)
  - Mica (e.g., biotite $\text{K}(\text{Mg,Fe})_3\text{AlSi}_3\text{O}_{10}(\text{OH})$).
The Giant Planets: Jupiter and Saturn

- Like the Sun, they consist primarily of H and He. Thus they are approximately “solar” in composition (though not exactly).
- Small rocky core under a metallic H layer, overlain by molecular H, He layer (gradual transitions from liquid to gas).
The Icy Planets: Uranus and Neptune

- These also consist of largely of H and He, but they are enriched in C and N compared to Jupiter and Saturn.

- Their solid parts consist of ices of methane and nitrogen.
Primary materials

<table>
<thead>
<tr>
<th>Materials in the Solar Nebula</th>
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</thead>
<tbody>
<tr>
<td><strong>Metals</strong></td>
</tr>
<tr>
<td>Examples</td>
</tr>
<tr>
<td>Typical Condensation Temperature</td>
</tr>
</tbody>
</table>

- The primary building blocks of the planets are metals (\(\rho \sim 8 \text{ g/cm}^3\)), silicates (\(\rho \sim 3 \text{ g/cm}^3\)), ices (\(\rho \sim 1 \text{ g/cm}^3\)) and H & He (\(\rho \) low, but varies with pressure).
- The bulk densities and other important properties of the planets are a direct result of the fact that materials with higher condensation temperatures will be able to condense closer to the Sun when the planets were forming.
Minimum mass solar nebula

- Consider the minimum amount of material that could have been in the solar nebula: this is the *minimum mass solar nebula*
  - take the solids from each planet
  - add volatile elements to each planet to augment them to solar composition (this represents the gas, which has dissipated)
  - spread each planet into an annulus reaching halfway to the next planet
  - smooth the resulting surface density $\Sigma$

- cylindrical coordinates, $R, \theta, z$ ($r^2 = R^2 + z^2$)
- azimuthally symmetric
- total mass $\sim 0.02 \, M_{\text{sun}}$ (lower limit)

$$
\Sigma(R) \approx 3 \times 10^3 \left(\frac{1 \, \text{AU}}{R}\right)^{1.5} \, \text{g cm}^{-2}
$$

$$
\Sigma_m(R) \approx 15 \left(\frac{1 \, \text{AU}}{R}\right)^{1.5} \, \text{g cm}^{-2}
$$

0.5% heavy elements (metals)
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Asteroid Belt (eroded by resonances)

Wiedenschilling 1975
Hayashi 1981
Optical depth

- Light travelling through the disk will reach the other side with intensity $I$ given by $I = I_0 e^{-\tau}$ where $\tau$ is the optical depth.
- Assuming the heavy elements are all (at least initially) in 0.1$\mu$m grains with $\rho = 3$ g cm$^{-3}$, the optical depth $\tau \sim 4 \times 10^5$.
- So the protoplanetary disk is optically thick.

middle B ring is optically thick in spots
inner C ring is optically thin

Varying optical depth in Saturn’s rings
Magnetic fields near the Sun

- The young Sun had a stronger magnetic field than its current field
- Ionized gas near the Sun would have been strongly coupled to this field, transferring angular momentum to the disk and slowing the Sun’s rotation (*magnetic braking*).
- At the *corotation point* (orbital velocity of gas = star’s rotation), the gas infall is effectively stopped by magnetic forces in the PP disk’s plane.
  - This gas can be lifted by magnetic forces out of the disk, and either deposited on the Sun or ejected, either to fall back into the disk further out or leave in a *bipolar outflow*.
- Heavy elements processed through this region may have resulted in the formation of CAIs and chondrules in meteorites.

Near the proto-Sun (Shu et al 1994)
A young star:

Accretion controlled by magnetic fields: `funnel flows'
Herbig-Haro object HH30 (HST)

IRAS 20068+4051
IRAS 13208-6020
The Egg Nebula
Hubble Space Telescope image of IRAS 13208-6020
Scale height

- assuming the disk is isothermal at a given R implies that its vertical density profile goes like

\[ \rho(z) \propto e^{-\frac{\Phi}{kT}} \]

where \( \Phi = \text{grav. potential} \quad \rho = \text{density} \quad T = \text{temperature} \)

\[ \Phi = -\frac{G\mu m_p M_{\text{sun}}}{(R^2 + z^2)^{1/2}} \approx \text{constant} + \frac{G\mu m_p M_{\text{sun}}}{R^3} z^2 \]

where \( \mu m_p = \text{molar mass of the H}_2 \) (dominant gas)

So the overall density profile is

\[
\rho(R,z) = \rho_0(R) \exp\left[-\frac{z^2}{H_z^2(R)}\right]
\]

where \( \rho_0(R) \) is the radial density profile (can get from \( \Sigma(R) \)) and \( H_z \) is the \textit{scale height}

\[ H_z^2 = \frac{2kTR^3}{G\mu m_p M_{\text{sun}}} \]

\[ H_z \approx 0.067 \text{ AU} \left(\frac{T}{500 \text{K}}\right)^{1/2} \left(\frac{R}{1 \text{AU}}\right)^{3/2} \left(\frac{2}{\mu}\right)^{1/2} \]
Flared disk

\[ H_z^2 = \frac{2kTR^3}{G\mu m_p M_{\text{sun}}} \]

We expect the disk to flare at larger R (though the R^{3/2} will be partly counteracted by lower T)

What we expect

Note: optically thick disks

What we see
Viscosity in the solar nebula

- Gas molecules orbit on roughly Keplerian orbits, and thus those closer to the proto-Sun move faster.
- Collisions tend to speed up molecules that are on the outside, and slow down those on the inside.
- This drives the disk to expand, with gas closer in falling on the proto-Sun.
- The mean free path between collisions $l_{fp} \sim 10$ cm.
- The resulting molecular viscosity $\nu = l_{fp} c$ is too small to cause any appreciable evolution of the disk ($c=$ sound speed).
  - This might imply that little gas from the disk ever makes it into the central star.
Viscosity in the solar nebula

• However, turbulent viscosity (mass and angular momentum transport by turbulence) is expected to be much more important.
• The *alpha parameter* $\sim 10^{-4} - 10^{-2}$ in the protoplanetary disk, which implies significant viscous evolution.
• The alpha parameter is an approximation that is also used in the study of other disks (around black holes, for example).

\[
l_{fp} = \frac{\mu m_p}{\pi r^2 \rho}
\]

\[
\nu_v \sim l_{fp} c
\]

\[
\nu_T = \frac{2}{3} \alpha_v c H_z
\]
Stability

• If the thermal velocity of the gas (or equivalently the sound speed $c$) is small, the collective gravity of a parcel of gas can overwhelm the tendency of its internal pressure and Keplerian shear to puff it up/stretch it out.

• Gravitational stability requires that Toomre’s parameter $Q_T > 1$

$$Q_T = \frac{c \Omega}{\pi G \Sigma}$$

$$\Omega = n = \left( \frac{GM_{\text{sun}}}{R^3} \right)^{1/2} \approx 2\pi \left( \frac{R}{1 \text{ AU}} \right)^{-3/2} \text{ yr}^{-1}$$

$$c = \left( \frac{\gamma k T}{\mu m_p} \right)^{1/2} \approx 1.7 \text{ km/s} \left( \frac{T}{500 \text{ K}} \right)^{1/2} \left( \frac{\mu}{2} \right)^{1/2}$$

• For standard parameters at 1 AU, $Q_T = 170$: the minimum solar nebula is very stable! This is a problem for the idea that planets collapse directly from the nebula (the direct collapse model) but as we will see, not for collapse of the solid component of the disk (core-accretion).
Direct collapse model

- Consider increasing the mass of the disk, to say 10x the minimum mass, and lowering its temperature, to say 100K.
- For these parameters, $Q = 2.4(T/100 \text{ K})^{1/2}$ roughly independent of radius.
  - So gravitational instability is just possible for extreme parameters;
  - direct collapse just might be invoked for Jupiter and Saturn and extrasolar gas giants, but not Uranus, Neptune, terrestrial planets (owing to their non-solar composition)
Direct collapse model

• ‘Alan Boss model’ (Science 1997)
• Pros:
  – Quite natural to form gravitationally unstable disks *if* there is enough mass or sound speed is low enough
  – Avoids the problem of dust agglomeration & meter-size barrier (later)
  – Planets forms quickly: no time scale problem (later)
• Problems:
  – Can disks get so very unstable? Gravitational spiral waves quickly lower surface density to marginal stability
  – Why do we have Earth-like planets?
• Simulation of a marginally unstable (Q=1.4) protoplanetary disk from 120-350 years (Mayor et al 2002). Several Jupiter mass planets are formed. Note Q=1.75 disk does not fragment.
Chemistry of Planet Formation

- The Sun constitutes over 99% of the mass of the Solar System. Therefore the solar nebula had an (elemental) solar composition.

- The solar nebula was composed of the interstellar medium, which evolved chemically during planet formation.

- The giant planets must have formed early, before the gas of the nebula dissipated, because they are rich in volatiles (H, He).

- Icy planets form more slowly because of lower densities in outer part of nebula.

- The inner planets form after the gas had largely dissipated from the inner solar system.
  - The terrestrial planets are depleted not only in gaseous elements such as H, He, C, and N, but in “moderately volatile” elements as well. These include the alkalis (Na, K, Rb, Cs) and elements such as sulfur, lead, and indium.

- Comets and chondritic meteorites were relatively unprocessed during planet formation.
Chemistry in the disk

• The raw material that forms the PP disk is heated by shocks as it falls onto the disk, and may be heated again later by further infalling material or proximity to the Sun.

• Where the temperatures exceed 2000 K, all elements are in the gas phase and the time scale to reach equilibrium is short enough that thermodynamic equilibrium can be assumed.

• At cooler temperatures, reaction rates slow and the competition between chemical reactions and condensation into the solid phase becomes complex to model.

• eg. in the solar nebula, CO is thermodynamically favored at T > 700 K and CH_4 at lower T. For nitrogen, N_2 is favored at T>300 K and NH_3 at lower T.

• however N_2 and CO are observed in the spectra of Pluto and Triton, implying that they cooled before they equilibrated chemically.
## Condensation sequence

<table>
<thead>
<tr>
<th>Temp (K)</th>
<th>Element</th>
<th>Form of Condensate</th>
<th>Comments</th>
</tr>
</thead>
<tbody>
<tr>
<td>2000</td>
<td>None</td>
<td>------</td>
<td>Gaseous Nebula</td>
</tr>
<tr>
<td>1600</td>
<td>Al, Ti, Ca</td>
<td>Oxides (Al₂O₃, CaO)</td>
<td>CAIs, etc</td>
</tr>
<tr>
<td>1400</td>
<td>Fe, Ni</td>
<td>Nickel-Iron grains</td>
<td>Planetary cores, iron meteorites</td>
</tr>
<tr>
<td>1300</td>
<td>Si</td>
<td>Silicates and ferrosilicates minerals</td>
<td>First stony material, combined to form meteorites</td>
</tr>
<tr>
<td></td>
<td></td>
<td>[enstatite, MgSiO₃; pyroxene, CaMgSi₂O₆; olivine, (Mg,Fe)₂SiO₄] in form of microscopic grains</td>
<td></td>
</tr>
<tr>
<td>300</td>
<td>C</td>
<td>Carbonaceous Grains</td>
<td>Forms soft black, carbonaceous rocks</td>
</tr>
<tr>
<td>less than 300</td>
<td>H, N</td>
<td>Ice particles (water, H₂O; ammonia, NH₃; methane, CH₄)</td>
<td>Large amounts of ice</td>
</tr>
</tbody>
</table>
Gas clearing: “magic broom”

- There is no primordial gas left between the planets, which implies it must have been cleared away at some point.
- Thought to occur when the Sun was $10^6$-$10^7$ yr old: T Tauri phase or Herbig-Haro phase.
  - much stronger solar wind?
  - luminosity 20-30x current: photo-evaporation of the disk?
- We’ll see the timing of the gas loss is very important to the core-accretion model, but is as yet rather loosely constrained.
The planetesimal (Safronov) hypothesis

- Planets form by multi-stage process:
  1. as the disk cools, rock and ice grains condense out and settle to the midplane of the disk
  2. small solid bodies grow from the thin dust layer to form km-sized bodies (*planetesimals*) - gas drag and solar gravity are dominant processes
  3. as particle mass increases, their gravity becomes important. Planetesimals collide and grow to become planets or planetary cores – gravitational scattering and solar gravity are dominant processes.
The planetesimal (Safronov) hypothesis

4. A few planetesimals grow large enough to dominate evolution: \textit{planetary embryos}

5. On much slower timescales, planetary embryos collide and grow into \textit{planetary cores} or \textit{proto-planets}

6. Cores of intermediate and giant planets accrete hydrogen/helium gas envelopes
   - requires growth by 45 orders of magnitude in mass through \approx 6 different physical processes!
Growth of solid bodies

• Chondritic meteorites are dated to 4.56 Gyr and formed within 20 Myr
• Evidence that they contained $^{26}$Al (half-life 0.72 Myr) indicates the solar system formed with a few Myr of an injection of freshly nucleosynthesized material (supernova or AGB stars)
• Alternatively, these radionuclides may have been produced by particle interactions near the active proto-Sun
• In any case, the gaseous atoms condense out of the nebula into small grains, with the exception of some interstellar grains, found in some meteorites, which seem to avoid any initial heating.
fractal dust aggregates

• Depending on the conditions, a variety of dust forms are expected as the grains hit and stick to each other

Theoretical models (Wright, UCLA)

Interplanetary dust (Brownlee et al)
Chondrules:
• formed at 4.55 Gyr with spread of only $10^7$ yr
• heated to 1500-2000 K and cooled in hours or less (melted fractal grains?)
• impact melting? magnetic field reconnection? solar flares? lightning? shocks? outflows?
Formation of solid particles

- As dust condenses out (metals in inner solar system, ices in outer solar system), the growth rate $dr/dt$ (where $T < T_c$)

\[ \frac{dr}{dt} \sim \frac{\alpha \rho_g}{\mu_p^{1/2} \rho_p} \]

- Growth rate depends on $\alpha$ (mass fraction of particulates in nebula), $c = \text{sound speed}$, $\mu_p$ the particle mass (in amu), $\rho_g \sim 10^{-9} \text{ g cm}^{-3}$ is the gas density and $\rho_p$, the particle density $\sim 3 \text{ g cm}^{-3}$.

- Yields $dr/dt \sim 1 \text{ cm/yr}$ independent of size
Dust settling time scale

- Dust suffers drag from gas in the nebula and will settle to the mid-plane of the disk on a time scale that depends on a competition between the vertical component of the disk’s gravity and (Epstein) drag ($\Omega = n = \text{orbital angular vel}, r = \text{particle radius}$)

$$- \frac{d\Phi}{dz} = -m\left(\frac{GM_{\text{sun}}}{R^3}\right)z = m\Omega^2 z$$

$$F_{\text{Epstein}} = -\pi r^2 \rho_g c v_z$$

- The settling velocity $v_z$ and timescale $\tau$ are thus

$$v_z = \frac{dz}{dt} = -\frac{z}{\tau}$$

$$\tau \approx \frac{\rho_g c}{\Omega^2 r \rho_p} \approx \frac{\Sigma}{\Omega r \rho_p} \approx \frac{160 \text{ yr}}{(r/1\text{ cm})}$$
Formation of solid particles

• So particles grow at ~ 1 cm/yr (independent of size) and settle to the midplane at a rate ~ $10^2$ yr/(r/1 cm)

• Therefore particles grow to a size of 10 cm as they settle to the mid-plane of the disk in about 10 years

• Problem: where do chondrules come from?
Formation of solid particles

• The gas disk is partly held up by its own internal pressure $P$, which partly counteracts the Sun’s gravity

• or, the effective mass of the Sun is reduced, reducing the orbital speed of the gas.

• As a result, the gas moves slightly (~0.2%) slower than the Keplerian velocity that the solid particles move at (particles feel little pressure support).

• Particle thus feel a “headwind” of ~ 50 m s$^{-1}$ with an attendant drag force per unit mass of

$$F = -X(v - v_g) \frac{m}{m}$$

where $X = \text{constant depending on gas density, geometry of dust grain, etc}$
The meter-size hurdle

- Small particles are entrained by the wind, large particles (hundreds of meters) are unaffected and intermediate particles spiral-inwards.
- The inspiral rate $dR/dt$ is given by

$$\frac{dR}{dt} \approx R \frac{2X \Omega (\Omega - \Omega_g)}{X^2 + \Omega^2}$$

- Maximum inspiral rate occurs where $X$ is numerically equal to $\Omega$, and is equal to the headwind speed
- The inspiral time for 30 cm particles $\sim 100$ yrs!
- Thus larger (km) particles (planetesimals) must form in less than 100 yrs or their constituents end up in the Sun
- This is called the meter-size barrier
There are two competing mechanisms for jumping the meter-size hurdle:

1. Gravitational instability (the Goldreich-Ward mechanism):
   - as solids settle to the midplane the particulate disk becomes gravitationally unstable when Toomre’s parameter
     \[ Q = \frac{c}{\Omega} \frac{\Sigma}{\pi G} < 1 \]
   - Here \( c, \Sigma \) are velocity dispersion and surface density of particles. If solid mass fraction is 0.5%, \( \Sigma = 15 \text{ g cm}^{-2} \) (minimum mass nebula) at 1 AU, instability requires \( c < 0.15 \text{ m s}^{-1} \) or thickness \( h = \frac{c}{\Omega} < 800 \text{ km} \)
   - For \( Q << 1 \) all wavelengths \( < \lambda_c = \frac{4\pi^2G\Sigma}{\Omega^2} = 1 \times 10^7 \text{ m} \) are unstable. Maximum unstable mass is \( M_c = \pi \Sigma (\lambda/2)^2 = 10^{16} \text{ kg} \), corresponding to final planetesimal radius of 10 km
   - Note: does not include “headwind” drag, which creates turbulent velocities that raise \( Q \sim 100 \). However, in the midplane where solid particle densities are high, they may overwhelm the gas and drag it along at Keplerian velocities, removing the headwind.
Jumping the meter-size hurdle

2. Sticky collisions

- particle velocities are turbulent \((v \sim 50 \text{ m s}^{-1})\) but collisions lead to sticking
- characteristic growth time \(r_{\rho_p} / \Sigma_p \Omega \sim 3 \text{ yr}\)
- but:
  - rocks don’t stick when they collide!
  - icy bodies fracture at these high speeds
  - largest inclusions in meteorites are a few cm
Formation of planetesimals

• once the meter hurdle is jumped, gas drag becomes unimportant

• gravitational interactions between planetesimals (bodies larger than 1 km) begin to be important:
  – enhance collision rate through gravitational focusing
  – fragments formed in destructive collisions are re-accreted
  – excite eccentricities and inclinations and thus pump up encounter velocities
Runaway growth

- When the relative orbital velocity of planetesimals is larger than the escape velocity at the surface of a planetesimal, mass growth goes like \((r = \text{planetesimal radius}) \ r^2\)
- If the relative velocities are smaller than the escape velocity, growth goes like \(r^4\) as gravity becomes more effective
- In the latter situation, a growing planetary embryo that ends up larger than its neighbours (possibly by chance) will rapidly increase its size and growth rate in what’s termed runaway growth

Simulation of runaway growth (Kokuba & Ida)
Oligarchic growth

- As its mass grows, an embryo begins to scatter planetesimals, increasing their relative velocities and slowing growth into the slower and more orderly oligarchic growth mode, where its own scattering (self-stirring) dominates the velocity dispersion of nearby planetesimals, rather than the disk’s inherent velocity dispersion.

- Occurs when the embryo reaches $10^{-6}$ to $10^{-5}$ Earth masses ($\sim 100$ km)
Isolation mass

- If the relative velocities are small, this implies the eccentricities of the particles are small, and so relatively few particles cross the orbit of the embryo.
- Embryos can reach material within a few Hill radii, which defines its (growing but slowly $\sim m^{1/3}$) feeding zone.
- As it depletes the material in its feeding zone, its growth slows to its isolation mass as embryos on near circular orbits settle in, separated by $\sim 5-10$ Hill radii.

$$R_H = \left(\frac{m}{3M_{Sun}}\right)^{1/3} r$$
Isolation mass

\[ M_{\text{iso}} = \left( \frac{\Sigma(t = 0)}{B} \right)^{1/3} \]

with

\[ B = \frac{3^{1/3} M_{\text{Sun}}^{1/3}}{2\pi b R^2} \]

• where \( b \) is the separation between embryos, in units of Hill radii
• Near the Earth’s orbit, the isolation mass would have been \(~ 6\) lunar masses, near Jupiter, \(~ 1\) Earth mass
• The embryo may be able to accrete more mass, but it depends on something else (eg. an already formed giant planet?) scattering material in.
• Clearly we have not reached the final planetary stage: how do we get these embryos, on near-circular, well-separated orbits to continue to accrete into larger bodies?
Dynamical friction

For first ‘half’ of growth one has:

$$\Sigma_M < \Sigma_m$$

In that case the planetesimal swarm dominates embryo by mass. *Dynamical friction* between planet and the swarm creates a drag on the embryo:

Dynamical friction:

Stirs up planetesimals (= creates ‘heat’ like friction) while damping the orbit of the embryo.

Dynamical friction circularizes the orbit of the embryo.
Final accretion stage

- The smaller planetesimals effectively acted as a gas, and drag from them had kept the growing embryos on circular orbits.
- However, as they are swept up, their influence diminishes and direct gravitational perturbations between embryos become important.
- Eccentricities of nearby embryos increase and they begin to cross orbits and accrete.
- In ~ $10^8$ years (in the inner solar system), the process is largely complete.

Chambers 2001
Terrestrial atmospheres

- The terrestrial planets are too small to hold significant amounts of H and He for long times but would have had primary atmospheres including these as well as other volatiles released but heating during accretion.
- Jean’s escape (thermal escape of light elements from a low mass planet) and impact erosion (ejection of gas by accretion impacts) would have affected the atmospheres in ways as yet not understood completely.
- In any case, the current atmospheres of the terrestrial planets are not primordial and are heavily altered, though they may contain some primordial tracers (noble gases).
Formation of the giant planets

• We have already discussed the direct collapse (gas-instability in deP & L) model
• The more accepted model is the core-accretion model (core-instability in deP&L)
• In this scenario, the giant planets form initially from a solid core as do the terrestrial planets but reach a larger mass owing to the ability of water ice to condense at larger distances.
• Once the proto-planet reaches 10-20 Earth masses, it can hold H and He gas from the nebula and rapidly increases in mass (recall the nebula is ~ 99% gas)
Runaway accretion

- Once beyond the ~10 $M_E$ threshold, gas begins to accumulate, slowly at first as the high temperatures following solid accretion create something of a pressure barrier.

- As the planet cools, gas begins to accumulate faster in *runaway gas accretion*.

- As the gas is expected to be removed on $10^6$-$10^7$ yr time scales, there is barely enough time for Jupiter and Saturn to form, a perceived difficulty with this model.

- The fact that Uranus and Neptune (14 and 17 $M_E$) contain relatively little gas indicates that they reached the runaway accretion threshold just as the gas was being cleared (possibly due to the lower density of solids at larger distances).
Type 1 Planetary migration

- occurs when a proto-planet does not yet have enough mass to open a gap in its gas disk
- It does generate a spiral arm in the disk, the mass of which is usually larger on the outside, resulting in a transfer of angular momentum that causes the body to drift inwards
- Lifetimes are calculated at $\sim 10^5(M_p/M_{\text{Earth}})^{-1}$ yr: this is an unsolved problem.

http://www.futura-sciences.com/img/planet-f1.gif
Type 2 Planetary migration

- occurs when the planet is large enough to open a gap in the disk, and also usually results in inward migration
- Migration is independent of mass, with time scales of $\sim 10^6$ yr
- Still a problem in our solar system perhaps, but may explain hot Jupiters in extrasolar planetary systems

http://www.astro.su.se/~pawel/planets/movies.html
Planetesimal-driven migration

- Once the gas is gone, the planets (outer in particular) continue to interact with the remaining planetesimals.
- Gravitational scattering is anisotropic, resulting in net angular momentum transfer.
- Jupiter probably moves in slightly, with the other giant planets migrating outwards.
- This causes mean-motion resonances (e.g., of Neptune) to migrate outwards, ‘sculpting’ the Kuiper Belt.
- Migrating resonances destabilize much of the remaining outer planetesimal disk, but some (e.g., 2:3, where Pluto and other Plutinos’s live) sweep up bodies.

Simulation of 4 planets and 10000 planetesimals (Morbidelli)
Asteroid belt

Questions:

- The mass of the asteroid belt is ~ \(10^{-3} \, M_{\oplus}\), much less than we’d expect from the smooth \(\Sigma(r) \sim r^{-3/2}\) for the initial expected distribution.
- Asteroids are also on more eccentric and inclined orbits, and have a variety of compositions (determined from spectra).

Answers(?):

- Jupiter is likely responsible for the lack of a planet in the AB, as well as its current sparseness and perturbed orbits.
- Composition differences may result from differentiation of some asteroids, temperature differences in the solar nebula, later “gardening” by impacts and chemical evolution through irradiation.
Formation of Kuiper belt and Oort cloud

1. Protoplanetary disc
2. Planetesimals form
3. Planets form
4. Interplanetary region cleared
5. Kuiper belt eroded
6. Inner and Outer Oort cloud form

Brett Gladman Science 2005
Outer planet satellites

- The outer planets have inner *regular satellite systems*, generally large moons on nearly-circular orbits in the plane of the planets' equator, and outer *irregular satellite systems*, smaller bodies on eccentric, highly inclined orbits.
Satellite systems

- Regular satellite systems probably formed from circum-planetary gas disks, similar to the solar nebula.
- Irregular satellites are thought to be captured planetesimals.
Terrestrial satellite systems

- The terrestrial planets are not massive enough to have had early circumplanetary gas disks
- Not surprisingly, they have few satellites
  - Mars has two small asteroid-like bodies probably captured from the asteroid belt
  - But what about Earth’s large Moon?
The Moon: key observations

- No other planet/dwarf planet has such a large moon relative to its size (except Pluto).
- The Moon has only a very small iron core (unlike Earth).
- The Moon has a bulk density about the same as the Earth’s mantle (suggests compositional similarity).
- Highly depleted in highly volatile elements (gaseous elements); depleted in moderately volatile elements.
- Has identical oxygen isotope composition to the Earth.
- Bottom line: Earth and Moon probably have a shared history but compositional difference make it unlikely they just accreted quietly next to each other.
Giant Impact Hypothesis

- Body about 1/10 the mass of the Earth (Mars-sized) struck the Earth after it is half or more accreted, and after the Earth’s core had at least partially formed.
- Material, mainly from silicate mantle, is blasted into orbit around the Earth, eventually accreting to form the Moon.
- Hence Moon has little iron, has a mantle-like composition and few volatiles.

Alternate idea: capture of small protoplanet
Problems: 1) straight capture is unlikely 2) lunar composition is not consistent with isolated formation anywhere in solar system.
Simulation of the giant impact

blue=iron, red = mantle. Note: fraction of mantle material in orbit later forms Moon
Simulation of the giant impact
The Nice model

- Named after the city in France (pronounced *niece*), this is a model for the late stage evolution of the planetary system.
- The basic idea is that the outer planets formed more closely to the Sun and subsequently underwent a rather violent transition to their current state.
The Nice model

• Initially, the outer planets extend no further than 17 AU (Neptune is currently 30AU) from the Sun. Jupiter and Saturn are near (but not in) 2:1 resonance

• Interactions between the planets and remaining planetesimals eventually push Jupiter and Saturn into 2:1 resonance, which causes a violent rearrangement of the Solar System to something like our current setup
The Nice model helps with

- Formation of U/N, since they form closer in
- Dispersal of Kuiper belt
- Late heavy bombardment (an increase in impact rates proposed to have happened ~800 Myr after the SS formed)
- There are no features of the solar system which have yet been found to be incompatible with the Nice model, and it does solve some outstanding problems, but it can probably never been proven rigorously.
The end