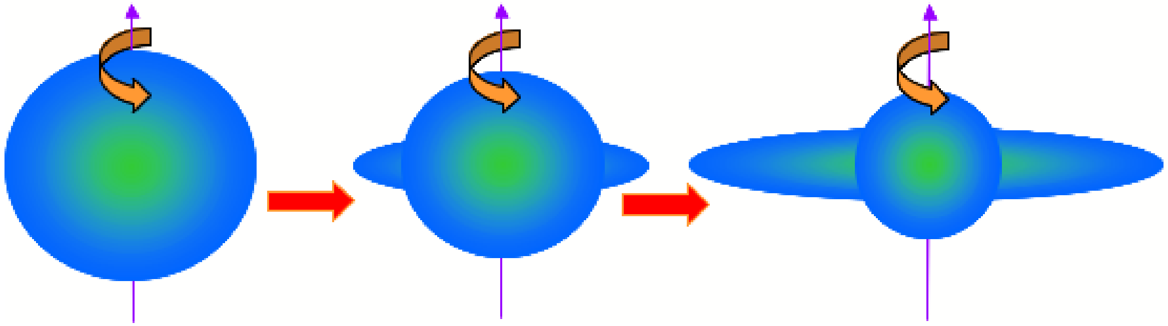
Planet Formation: Disk Formation and Evolution



Summary

In this Activity we will look at the formation and evolution of protoplanetary disks in which planets are formed. In particular we will discuss:

- some of the constraints from the Solar System that planet formation models must explain;
- the standard star formation paradigm and details of cloud collapse;
- disk formation, evolution, and dispersal;
- chemistry and dust grain condensation; and
- grain growth via collisions.

Introduction

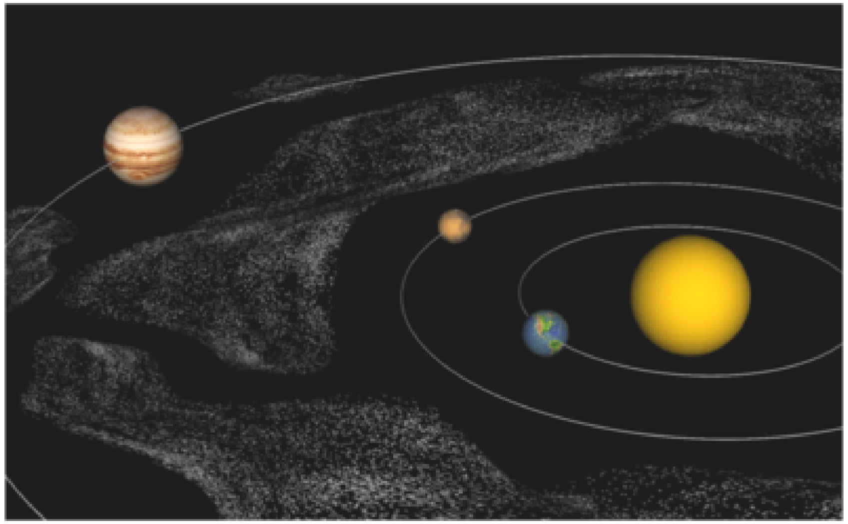
The formation of the planets and the Solar System is one of the fundamental questions in modern astronomy. It has been a question asked throughout the ages since it is really a question about our own origin. We cannot hope to fully understand how life evolved on Earth if we do not know how the Earth itself formed.

Throughout this course we have found many clues to help us build a general picture of the planet formation process, as well as a number of constraints that any planet formation model must be able to address. As well as planets in our own Solar System, we now know of hundreds of planets orbiting other stars in the neighbourhood of our Sun. Assuming our Solar System in not cosmically special, a general planet formation paradigm should be able to understand the formation of the extrasolar planets as well.

In this Activity we will start by looking at some of the constraints from our Solar System, and then briefly review the standard star formation paradigm, since planetary disks are a natural by-product of the star formation process. Once formed, protoplanetery disks go through a complex evolution with can result in the formation of planets, satellites, asteroids and comets before it dissipates within a few tens of millions of years. We will look at disk evolution, as well as how grains form and grow in these disks. In the next Activity we will look more closely at models of terrestrial and giant planet formation.

Constraints from the Solar System

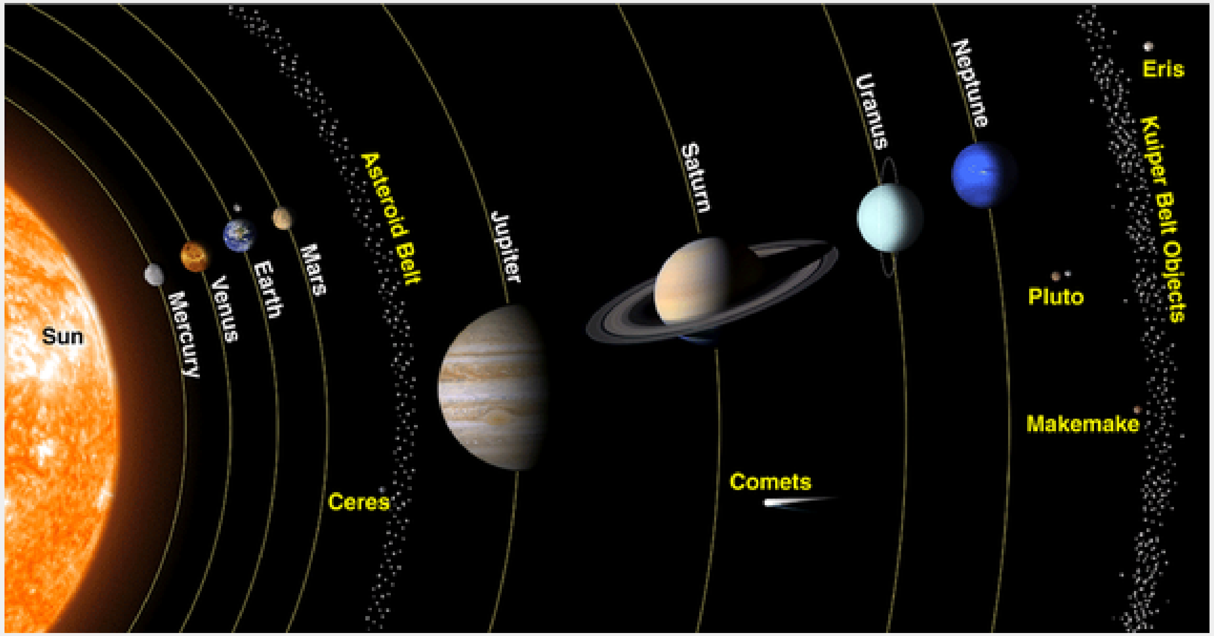
If we take stock of what is in the Solar System, we find a wide range of objects: the central hot sun, rocky planet in the inner Solar System, gas and ice giants in the outer Solar System, a mix of rocky and icy satellites, as well as a mix of rock and icy minor planets, and icy comets. Any model of planet formation needs to explain all of the objects, and a model for our Solar System needs to include an evolutionary scenario that explains the current state (location, composition, age, etc.) of all of these objects.



The geometry of the Solar System provides important clues, as well as constraints, on formation scenarios. The planets (major and minor) reside in the same plane and are relatively regularly space; they all orbit the Sun in the same direction; most planets spin in the same direction in which they orbit the Sun; and orbital periods increase with distance from the Sun. We also see a range of dynamical structures in the asteroid belt and the Kuiper belt and these smaller bodies are in general on more eccentric and inclined orbits than the planets.

In terms of the composition and bulk chemistry, we find the smaller rocky planets (Mercury, Venus, Earth, Mars) in the inner part of the Solar System, while the larger gas/ice giants (Jupiter, Saturn, Uranus, Neptune) reside farther out. The density of the planets decreases with heliocentric distance. There is an inner ring of small rocky bodies, i.e. the asteroid belt (which has a chemical gradient) and an outer belt of icy bodies, i.e. the Kuiper belt. In the centre of the Solar System is the hot Sun. There are clearly more refractory materials close to the Sun and more volatiles farther away.

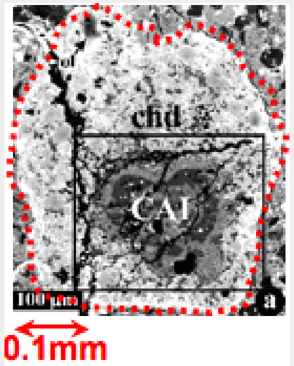
Small bodies are generally irregularly shaped and undifferentiated, while large bodies are spherical and differentiated. This tells us something about the history of internal heating of larger planetary bodies. The surfaces of planetary bodies, which can be crater and/or lava covered, also provide clues about their thermal history. Crater counts also provide information about impact rates in the early Solar System.



Credit: NASA

The refractory inclusions in the meteorite data tell us that the first solids in the Solar System formed about 4.568 Gyrs ago. The chondrules and differentiated planetesimals formed a few million years later. Lunar rocks are between 3–4.4 Gyrs, and the oldest Earth rocks are about 4 Gyrs old (though terrestrial grains with ages $\sim$4.4 Gyrs have been found).

We know that planets are formed around stars as a natural by-product of the star formation process. So before we move on to the details of planet formation, we need to start with the general theory of star formation.



First condensates: calcium aluminium inclusions.

Credit: © Sasha Krot

The standard star formation paradigm

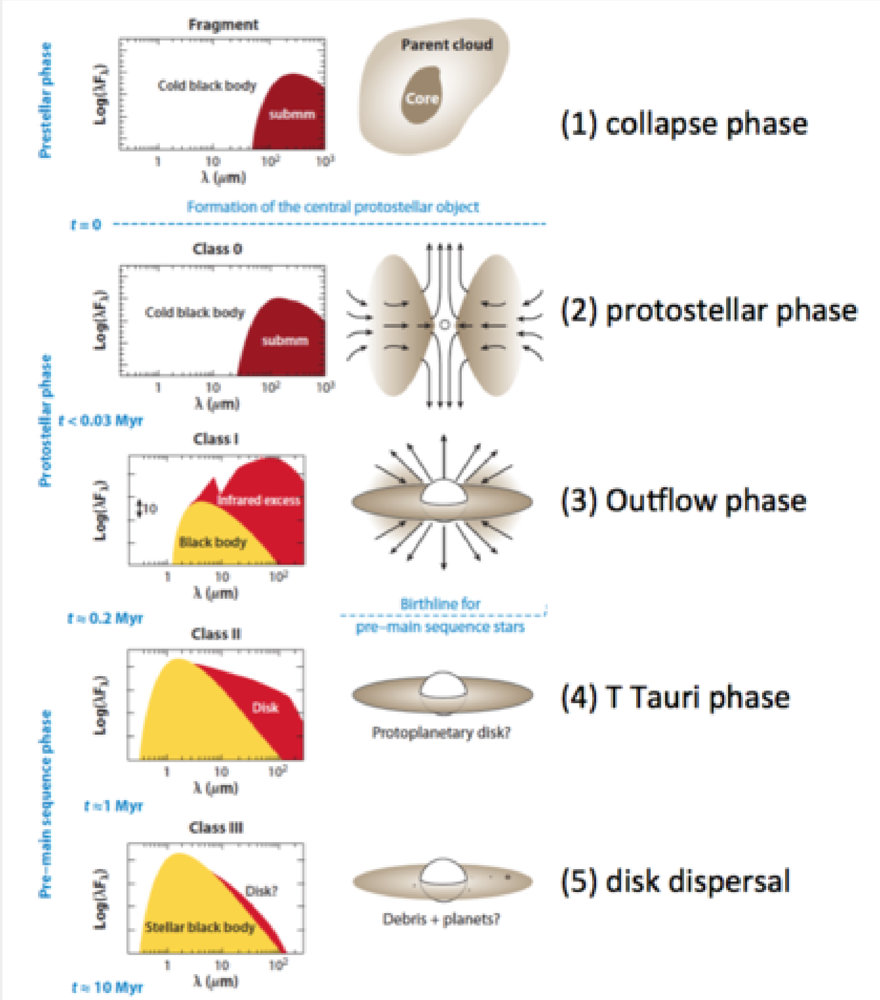
Stars form in molecular clouds that are found in the spiral arms of the Galaxy. Molecular clouds are cold, dense clouds of gas and dust, and stars form within them when they contract due to gravity.

The Milky Way contains a large number of molecular clouds, ranging from giant systems with masses in the range of 10^5–10^6 $M_\odot$ and sizes in the order of 50–100 pc, down to small cores with masses as low as a few $M_\odot$ and a few pc in size. The small cores are usually embedded in larger complexes. Most molecular clouds have temperatures between 10–50 K and as their name suggests contain a wealth of molecules. Densities range from 10^3 particles per cm^3 in molecular clouds to 10^5 particles per cm^3 in dense cores.

In spiral galaxies, stars form in the arms which contain dense regions of gas and dust.
Credit: NASA and The Hubble Heritage Team (STScI/AURA)

The generally accepted paradigm of low mass star formation (Shu et al. 1987) is as follows:

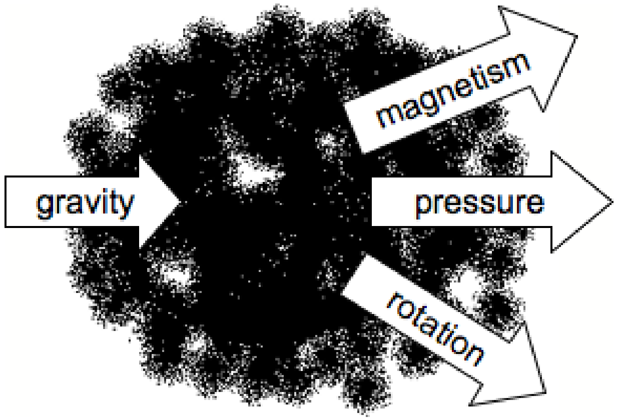
1. **collapse phase**: giant molecular clouds must contract to form molecular cores. This contraction requires *ambipolar diffusion* to first carry away the magnetic fields which help hold the cloud up;
2. **protostellar phase**: the rapid inside-out gravitational collapse of molecular cloud cores conserves angular momentum, producing a protostar surrounded by a disk and an optically thick infalling envelope;
3. **outflow phase**: a strong stellar wind breaks out at the rotational poles, reversing the infall and producing bipolar outflows. This phase seems to be intimately connected with the disk formation phase;
4. **T Tauri phase**: the newly formed star/disk system becomes optically visible and the protostar is identified as a T Tauri star;
5. **disk dispersal phase**: the final stage is the clearing of the disk, via photoevaporation and stellar winds.



To view a full-sized version of this diagram click here.

Cloud collapse

Giant molecular clouds are held up by **magnetic fields**, thermal gas **pressure** and **rotation**. If the clouds become massive enough, they can collapse due to **gravity**. We will briefly look at each of these support mechanisms.



The role of turbulence in star formation is rather complex: it creates overdensities which can initiate gravitational collapse and it can also counter the effects of gravity in these overdense regions (McKee & Ostriker 2007).

(i) Pressure support: Pressure is related to temperature via the gas *equation of state*. Therefore inside cold molecular clouds the **thermal pressure** will not provide much support against gravitational collapse. Some clouds, however, have observed supersonic motions[1] and thermal gas pressure which may contribute to their stability.

(ii) Rotational support: The gravitational force that is trying to make the cloud collapse also has to overcome **rotation**. The cloud's rotation produces an outward centrifugal force perpendicular to the rotation axis that keeps the cloud from collapsing. Observations of molecular clouds indicate that they rotate quite slowly, with rotational velocities equivalent to 0.3 km/s at a distance of 0.1 pc. This is similar to the thermal support, but many clouds rotate *slower* than this which suggests that rotation provides little support[2].

[1] Supersonic means faster than the speed of sound. The speed of sound in an ideal gas depends on the density and pressure.

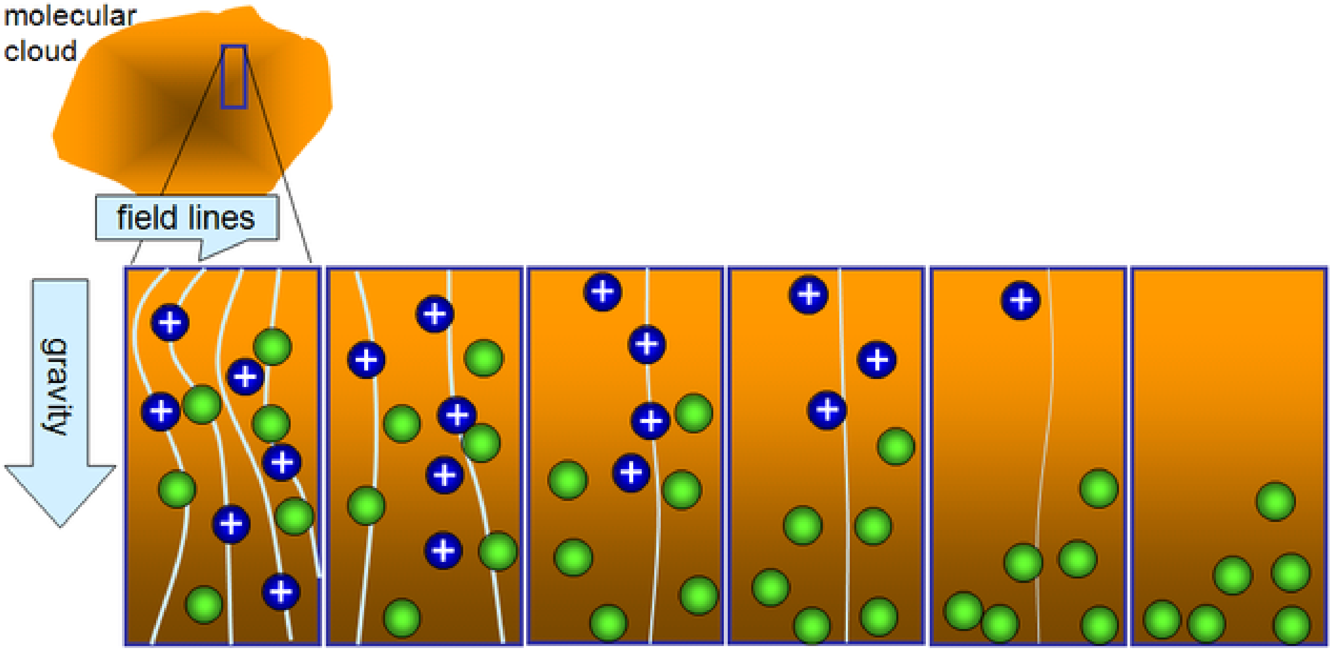
[2] While rotation may not be important in holding clouds up against local collapse, we will see that rotation has an important role to play in the star and planet formation process.

(iii) Magnetic support: **Magnetic fields** only affect charged particles in molecular clouds, making them follow field lines rather than obey gravity (or dragging the field along with them). The fields exert a magnetic force on the charged particles, which acts like a *magnetic pressure*. The magnetic pressure is given by:

$$P_{mag} = \frac{B^2}{8\pi}$$

It is difficult to measure the magnetic field strength in molecular clouds. In a study of 27 clouds Crutcher (1999) found field strengths B = 20–200 μG. This is generally thought to be enough to support molecular clouds against gravitational collapse and is about ten times greater than the thermal support.

In order for molecular clouds to collapse and commence the star formation process, they must rid themselves of this magnetic support. The process which removes magnetic pressure from molecular clouds is called ***ambipolar diffusion*** (Mestel & Spitzer 1956, Mouschovias 1976). In this process there is a very slow slippage between the ions and neutrals in the cloud, and the ions carry the magnetic field with them (whereas the neutrals are unaffected by the magnetic field). Thus the ions slowly diffuse away the magnetic field. Ambipolar diffusion is also known as “ion-neutral drift”.



Ion-neutral flows have been observed in molecular clouds (Greaves & Holland, 1999) supporting the theoretical idea of ambipolar diffusion. Once the magnetic support has been fully removed, the cloud can finally collapse due to gravity.

Core collapse

For a cloud in equilibrium, the ***Virial theorem*** states that the gravitational potential energy is twice the kinetic energy, $\mid E_{grav} \mid = 2E_{kin}$, where the kinetic energy mainly comes from thermal energy (unless the cloud is very turbulent or rapidly rotating).

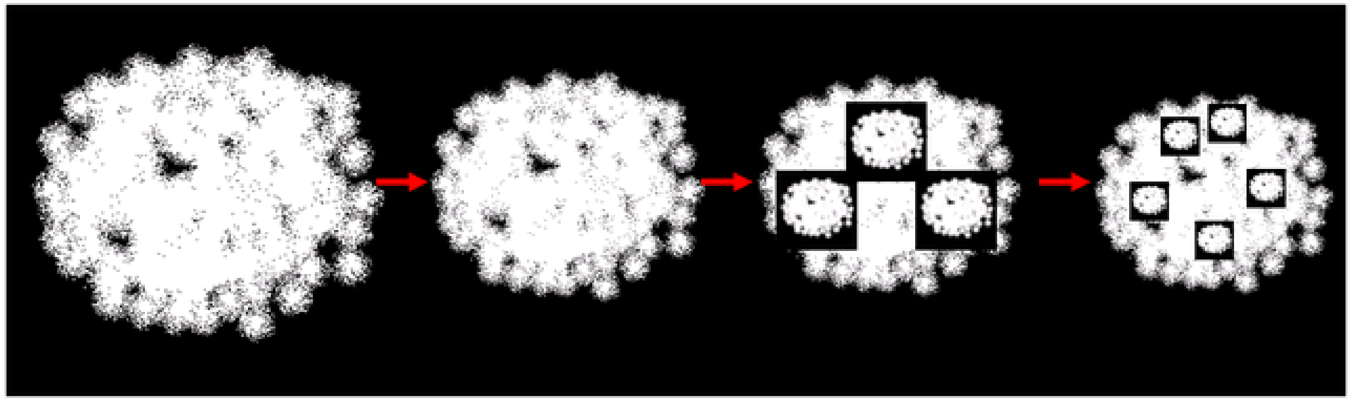
When $\mid E_{grav} \mid > 2E_{kin}$ the core can collapse. We can solve the Virial equation to find the minimum mass - called the ***Jeans mass*** - of the cloud that will lead to gravitational collapse (if supported by thermal pressure alone):

$$M_{Jeans} \propto \sqrt{\frac{T^3}{\rho}}$$

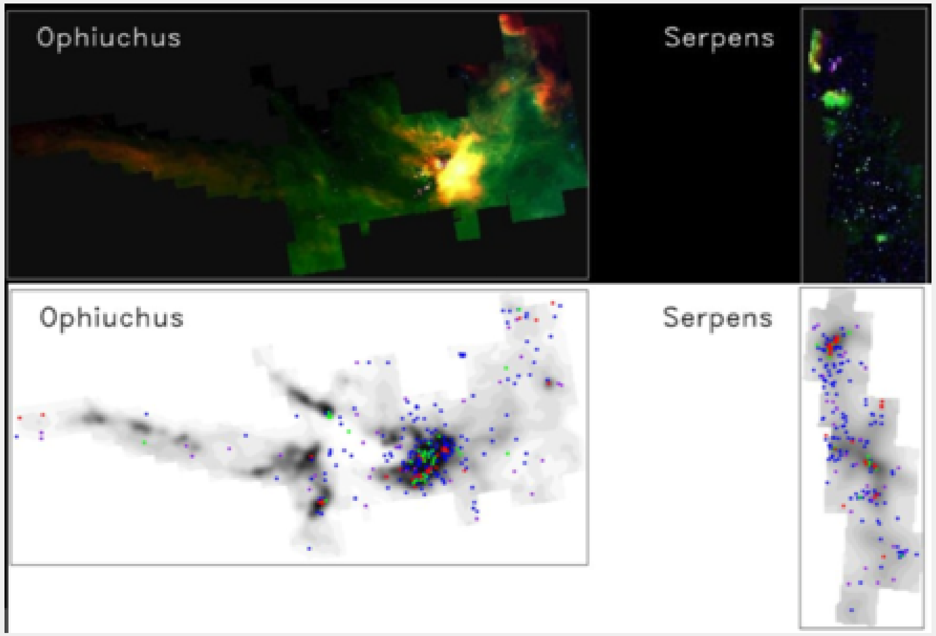
Note that the Jeans mass, M_{Jeans}, decreases as the density, ϱ, increases. Observations of cloud cores indicate that they are dense enough to collapse to a stellar mass object, but we do not see densities high enough that cores can collapse directly to produce Jupiter mass (i.e. planetary) objects.

Fragmentation

Let's assume that the Jeans criterion is satisfied and the cloud is unstable to gravitational collapse. As the cloud contracts, different regions within the cloud will satisfy the Jeans criterion individually and start to collapse themselves. Therefore many *smaller* objects are formed, and this process is known as **fragmentation**.



Note that the density of the cloud increases by many orders of magnitude during the free fall process, yet the temperature remains nearly constant. The Jeans criterion tells us that the Jeans mass must decrease as the collapse continues. Thus small regions of the collapsing cloud will satisfy the Jeans criterion locally, resulting in many small stellar cores. This is supported by observations of star forming regions, which show that stars generally form in clusters covering a range of sizes.



Spitzer images of the star forming regions Ophiuchus and Serpens from Evans et al. (2009)

Credit: Tyler Bourke, CfA

Protostar phase

As the molecular cloud core collapses, energy must be conserved and the gravitational potential energy is converted into thermal energy. While the dust in the core is optically thin, the core is transparent to infrared wavelengths, and the thermal energy gained by the collapse will be radiated away and the core stays cool. This phase of contraction is called **isothermal collapse** during which the temperature remains about constant.

The increase in density during the collapse, however, causes the optical depth to increase and eventually makes the cloud opaque to infrared, trapping the thermal energy. This is called the **adiabatic collapse** phase, when energy is not radiated away and the temperature increases as the core collapses. The trapped thermal energy heats the core, which builds up the internal pressure until hydrostatic equilibrium is reached (i.e., gravity balances pressure). As the temperature rises, the molecular hydrogen, H_2, dissociates into separate hydrogen atoms. The end product is a quasi-static protostellar core surrounded by an in-falling envelope. At this stage the core is officially called a **protostar**.

Protostars shining through the Eagle Nebula in the infrared.
Credit: M. McCaughrean "& M. Andersen, Astrophysical Institute Potsdam, ESO.

Once the central temperature reaches about 10^6 K, nuclear reactions begin in the core and convert deuterium (D or ^{2}H) into helium (He) via the reaction:

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

This nuclear energy heats the core and halts any further core collapse until the deuterium is exhausted. Then the protostar shrinks again and heats up until the temperature reaches about 10^7 K. Finally hydrogen fusion is ignited in the stellar core, and a star is born!

$$4\,{}^1H \rightarrow {}^4He + 2e^+ + 2\nu$$

An aside: some definitions

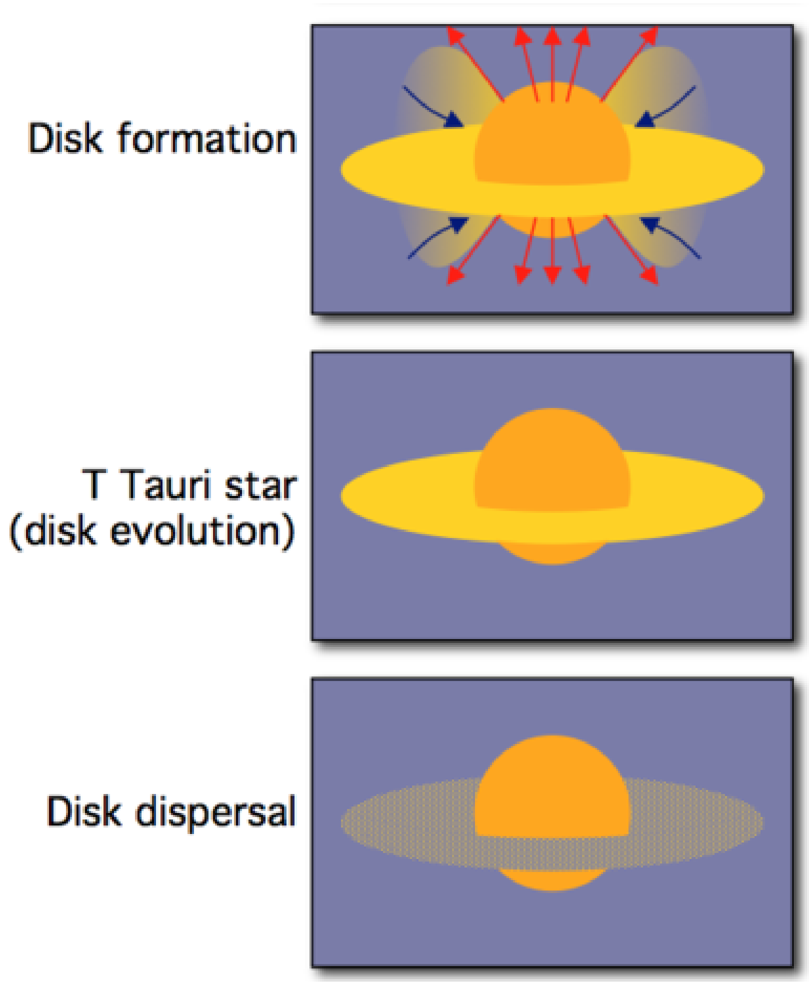
There are a wide variety of terms used when referring to young stars, including *protostar*, *pre-main sequence star*, *T Tauri star*, and *young stellar object*. In 1972, Steve Strom coined the term **young stellar object** (or YSO for short) in recognition of the circumstellar material that strongly affects and alters the appearance of young stars. The YSO can refer to T Tauri stars, Herbig Ae/Be stars, protostars, and pre-main sequence stars. So what are all these objects?

- **Pre-main sequence stars** (or PMS stars) are stars that are in the process of contracting onto the main sequence and are not yet burning hydrogen in their cores.
- A **protostar** is an embedded source that has not yet broken free from its parent cloud core (while PMS stars are revealed sources, meaning that they have broken free from their parent cloud core). Like PMS stars, protostars are not yet undergoing hydrogen fusion.
- Low mass ($M < 2\ M_{\odot}$) YSOs are called **T Tauri stars** (TTS).
- High mass (2–10 $M_{\odot}$) YSOs are called **Herbig Ae/Be star** (HAEBE).

Disk formation and evolution

During the gravitational collapse of a molecular cloud core, the gas must contract by a factor of about 10^6 in size to form a star. Due to conservation of angular momentum, the initial cloud rotation is enormously magnified, which results in the small central protostar being surrounding by a **large rotating disk**. It is in these disks that planets form.

In this section we will discuss the three main stages in the life of a protoplanetary disk: formation, viscous evolution, and dispersal.



An aside: some more definitions!

You will see various expressions used when reading about the disk out of which our Solar System formed.

- The **solar nebula** refers to the proto-Sun along with the surrounding disk, though it is often used (erroneously) in a way that means just the disk component.
- The disk itself is usually referred to as a **protoplanetary disk**.
- The **minimum mass solar nebula** is a theoretical disk that contains the minimal amount of mass (with the appropriate composition) to form the planets[3].

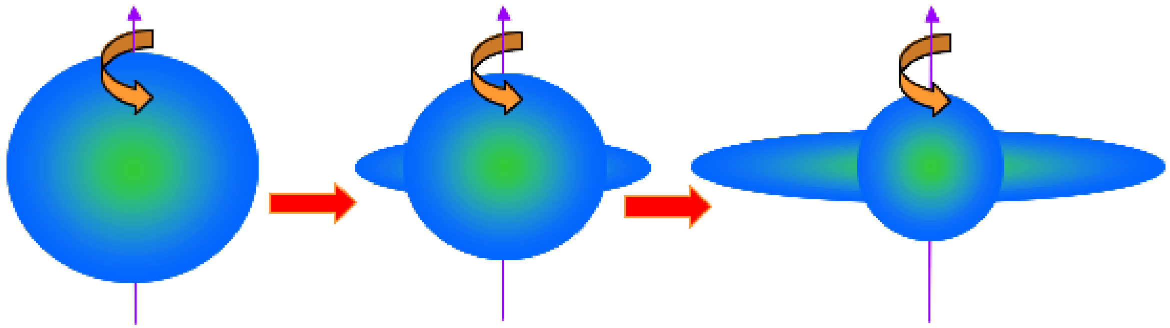
[3] The term "minimum mass solar nebula" was coined by Weidenschilling (1977) and technically defines a mass distribution. This mass distribution can be integrated within a specific radius to get the minimum mass, which is generally between 0.01–0.02 $M_\odot$.

Disk formation

Protoplanetary disks form as a natural consequence of cloud collapse due to angular momentum conservation. Angular momentum is given by

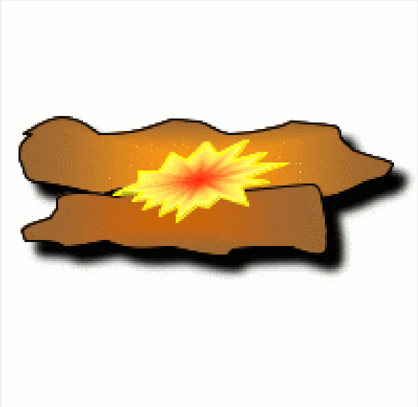
$$\overrightarrow{L} = m\overrightarrow{r} \times \overrightarrow{v}$$

So as the cloud contracts, *r* gets smaller thus *v* must increase and so the cloud spins faster. This rapid rotation creates large **centrifugal forces** - just like you feel in a car when you go too fast around a corner, and large sideways forces push you away from the corner. These centrifugal forces are greatest at the equator and the rotating contracting cloud starts to spread out and form a disk:



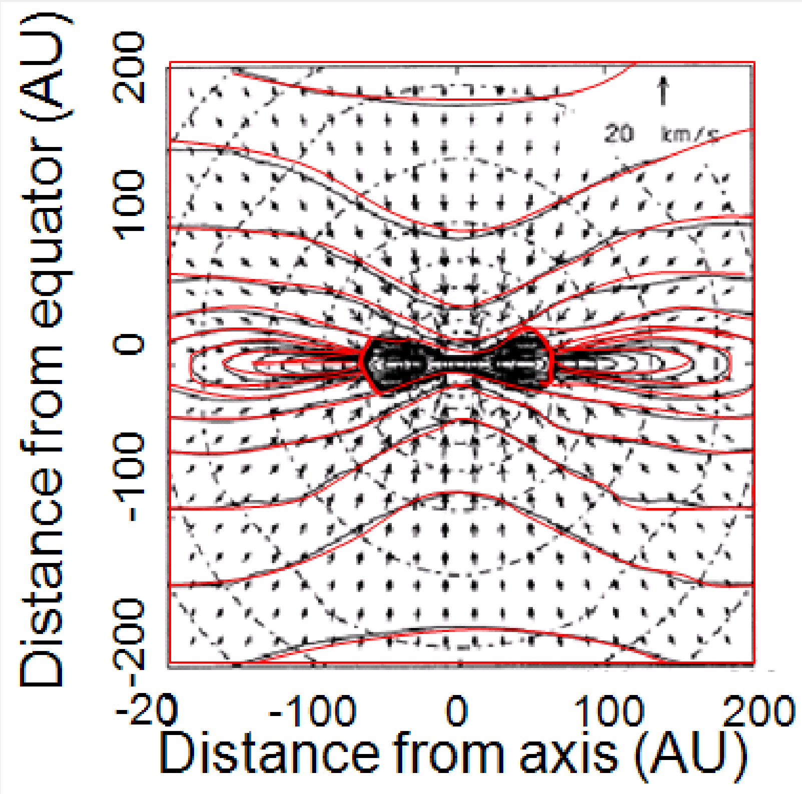
The infalling gas rapidly reaches supersonic velocities as it falls to the mid-plane, and passes through a shock front at the equator of the collapsing cloud. The shock dissipates the kinetic energy of motion perpendicular to the disk. If the shocked gas cools quickly, then the material will accrete onto the mid-plane and settle into a thin disk.

Evidence of accretion shock in protostellar disks was found in the system Lynds 1157 (Velusamy et al. 2002)



Shocked gas in the disk midplane due to infall.

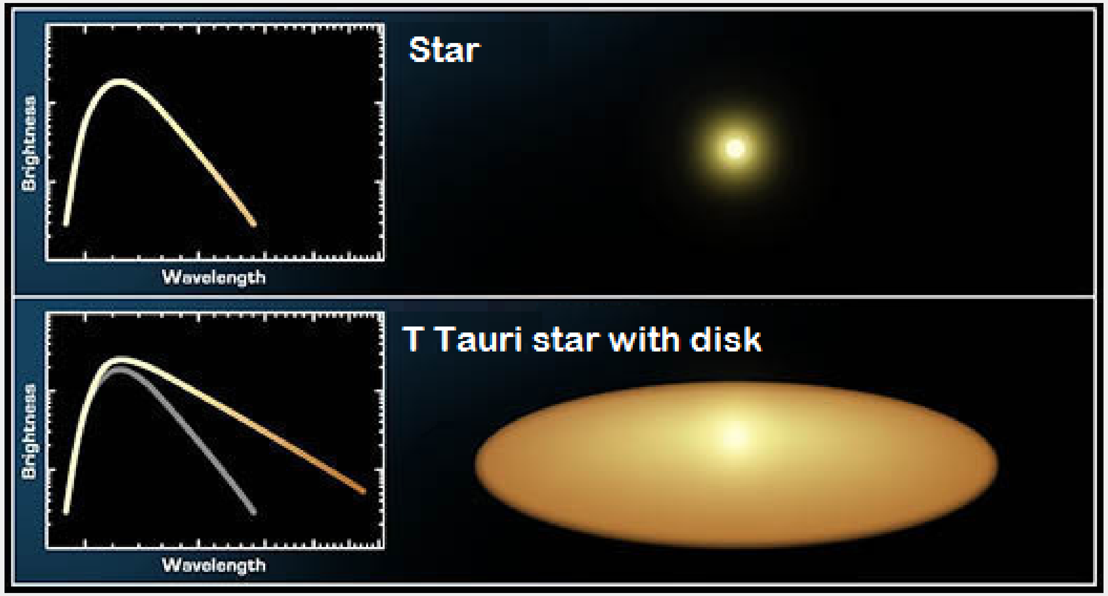
The disk will reach a quasi-equilibrium when all the forces balance: the inward gravitational forces towards the centre must balance the outward centrifugal force, and the gravitational force towards the disk centre must be balanced by the pressure gradient in the disk.



Results of a rotational cloud collapse simulation by Yorke et al. (1995). Equal density contour lines are shown in red.
Credit: H.W. Yorke, P. Bodenheimer, G. Laughlin (1995)

Disk evolution

Observations of young stellar objects and T Tauri stars in particular are generally done at sub-millimeter and infrared wavelengths. If we plot the spectral energy distribution of a T Tauri star, we find that there is excess infrared emission. This is due to the dust in the protoplanetary disk, which absorbs stellar radiation and re-radiates in the infrared. These disks are called “reprocessing”, “irradiated”, or “passive” disks.



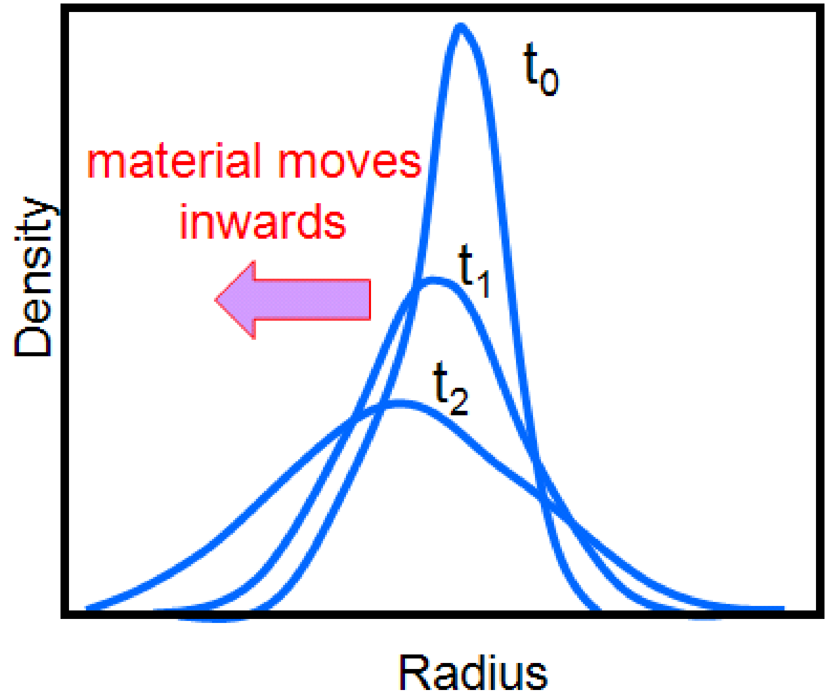
Excess infrared emission is seen in T Tauri stars due to their disks.

Credit: NASA/JPL-Caltech/T. Pyle (SSC)

Some disks, however, are *much* brighter than can be accounted for by this reprocessing, and therefore they must have some energy source of their own. These disks are called “active” or “accretion” disks since the source of energy comes from the accretion of material within the disk.

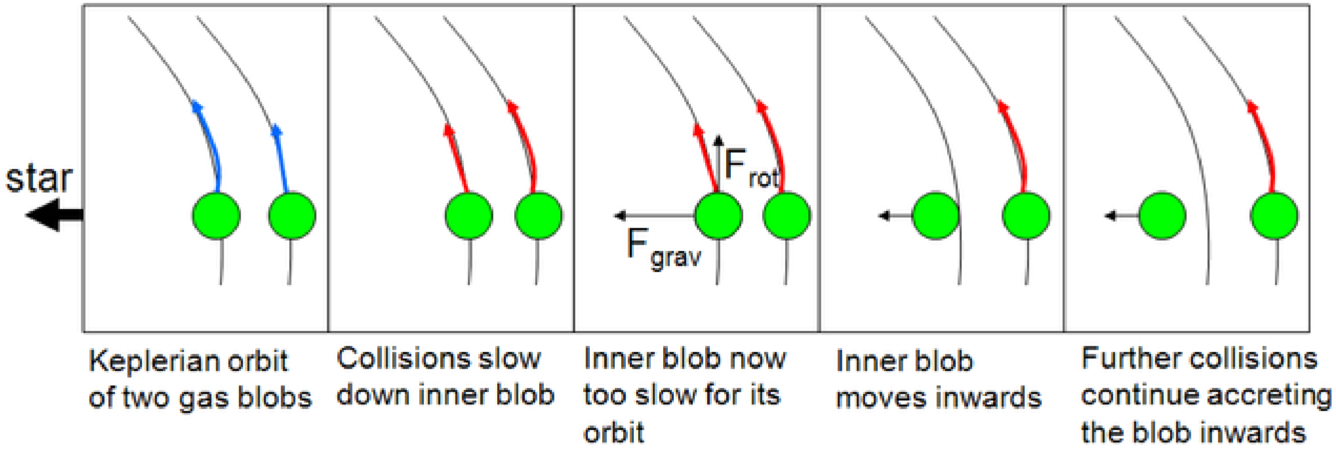
During the accretion phase angular momentum is transferred outwards and material moves inwards and accretes onto the protostar. The efficiency of the angular momentum transport determines the structure and evolution of the disk. Lynden-Bell & Pringle (1974) considered mass and angular momentum conservation to solve for the density as a function of radius and time. They showed that a viscous disk will spread radially, redistributing the density so that material moves inwards and angular momentum is transferred outwards.

Various mechanisms have been proposed for the transport of angular momentum and mass in accretion disks, including viscous torques, gravitational torques, and magnetic torques.



Viscous torques

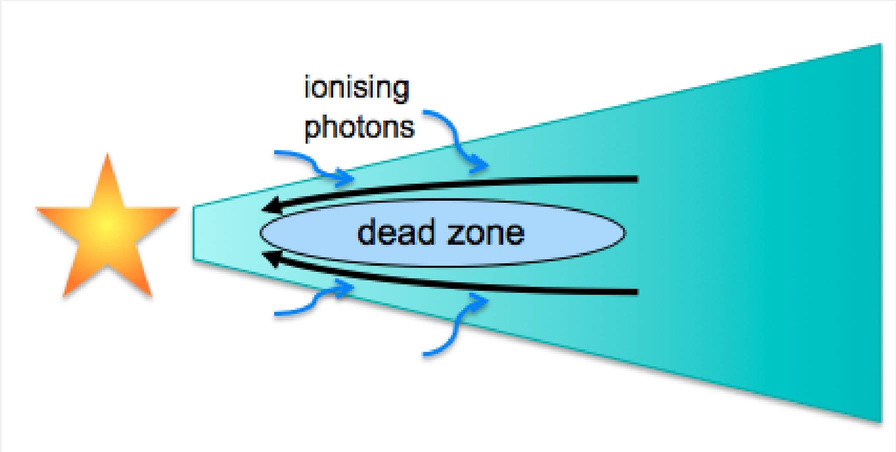
Gas in the disk follows approximately Keplerian orbits, so material in the inner part of the disk moves faster than material further out. Collisions between neighbouring blobs of gas speed up the outer blobs and slow down the inner blobs. The inner blob will now be moving too slow for its orbit, and so will fall inwards.



The reverse is true for the outer blob, which will move outwards.

Thermal motions of the gas produce *molecular viscosity*, but this has been found to be too small to result in any significant evolution over the disk lifetime. *Turbulent viscosity* is more likely to result in viscous disk evolution. Temperature gradients within the disk can cause convection, and convective eddies can induce radial mixing of disk material while transferring angular momentum in the process. The direction of the angular momentum transfer due to convection is debated.

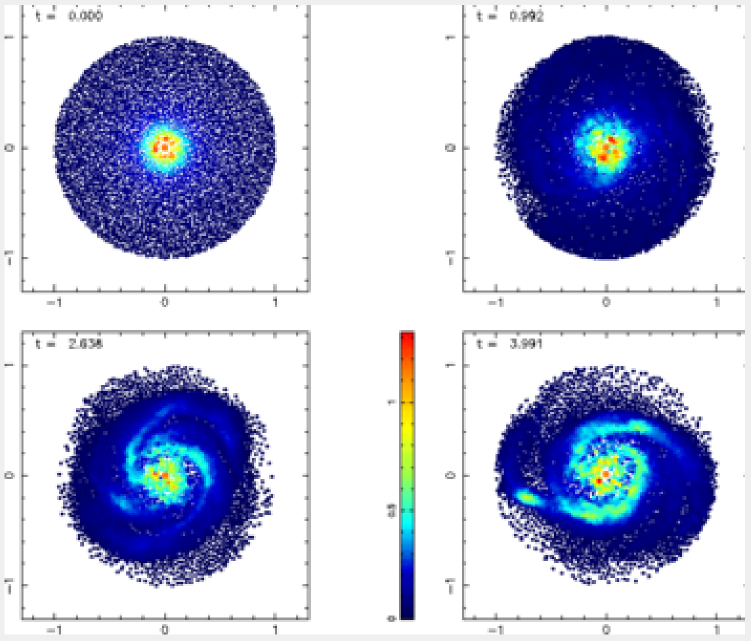
Magnetohydrodynamic (MDH) turbulence, on the other hand, is much more promising (Balbus & Hawley 1991). Shear flows in weakly magnetised disks can lead to very strong local instabilities, producing sustained turbulence which drives angular momentum outwards. Even a small amount of photon or cosmic ray ionisation in the disk will result in weak ionisation. Near the disk mid-plane, however, the disk is well shielded from ionising radiation and MHD turbulence becomes ineffective. These regions of very low viscosity are called *dead zones* (Gammie 1996) since accretion is inactive.



Dead zone within an accretion disk, with ionising photons cannot penetrate and viscosity is very low.

Gravitational torques

Soon after its formation, the disk is likely quite massive as it continues to accrete material from the infalling molecular cloud core. Massive disks are prone to both local and global gravitational instabilities. This can lead to the formation of spiral density waves which are very efficient at redistributing angular momentum (and material) through the disk until quasi-equilibrium is reached.



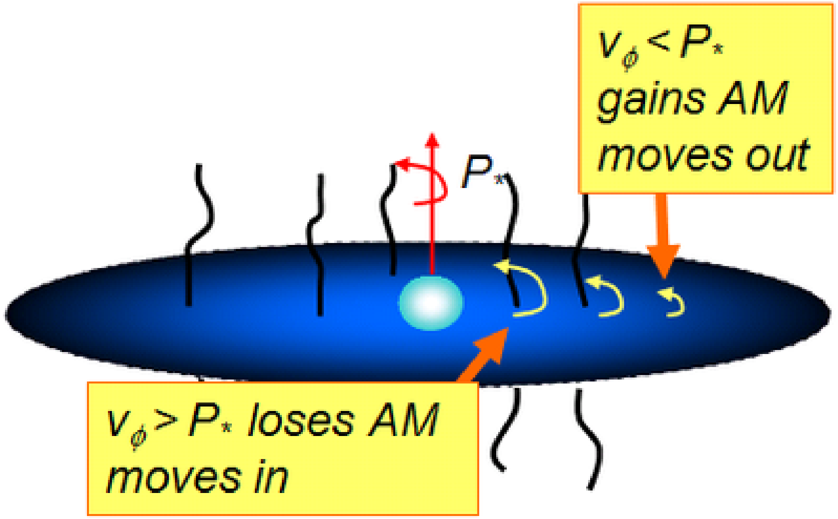
Simulation of massive disk with M_{disk} = 0.5 M_*.

Credit: Sarah Maddison (Swinburne University)

Magnetic torques

If the magnetic field lines from the protostar thread the disk, then the field tries to drive the disk to *corotation* - which means it tries to make the disk spin at the same rate as the protostar's rotation period, P_*. Disk material orbiting more rapidly than P_* will therefore lose angular momentum, and material orbiting slower than P_* will gain angular momentum.

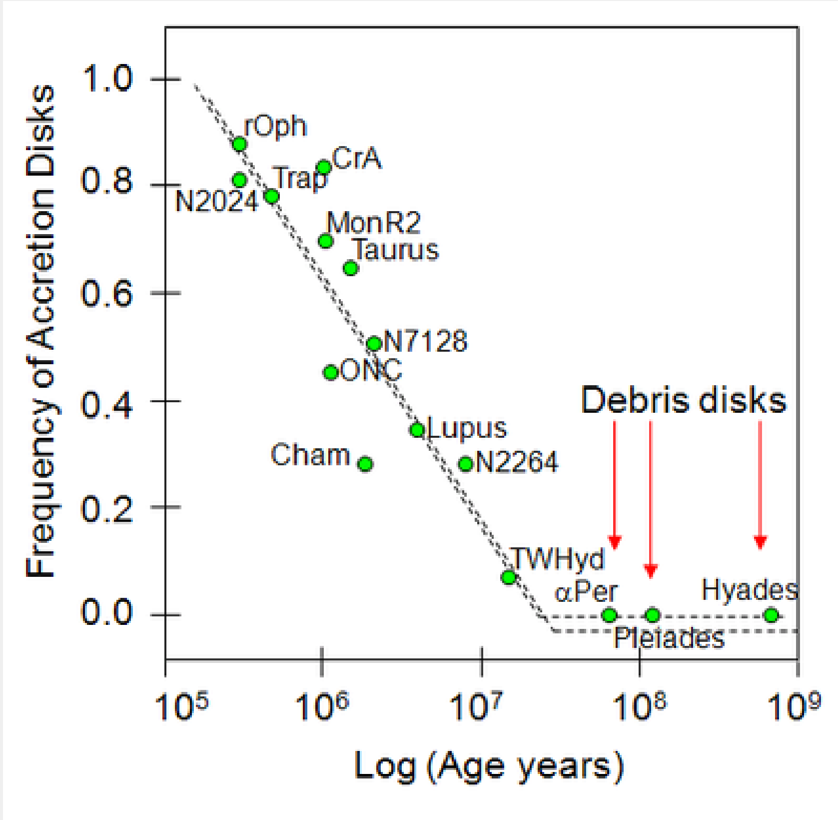
The hot inner disk will be ionised, and so the magnetic field couples the star to the disk. The rapidly rotating protostar will be slowed down by *magnetic breaking torques*, and angular momentum will be transferred outwards from the star to the disk.



Disk dispersal

The final stage of disk evolution is disk dispersal. The oldest observed (gas-rich) protostellar disks are about a few 10^6–10^7 years old (Haisch et al. 2001), so the disk must be cleared within this time frame.

Obviously, the timescale of disk clearing strongly constrains the planet formation process, because once the disk has dispersed, there is no more material available to make planets.



Disk fraction as a function of mean cluster age of various star forming regions. The disk fraction comes from IR excess observations.

Credit: Swinburne, from Haisch et al. (2001)

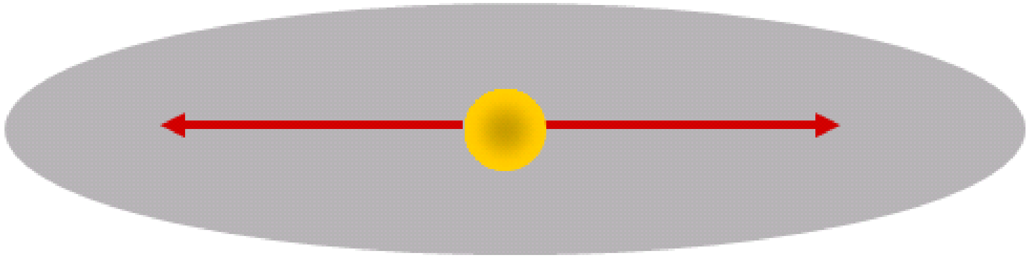
Mechanisms for removing disk material include accretion onto the protostar (as we've been discussing) as well as forming planets. Disk clearing is also assisted by the stellar wind and photoevaporation of the surface of the disk by the high energy radiation from the young protostar (or nearby massive stars). It is quite probable that all of these processes play a role.

Hollenbach & Adams (2004) suggest that photoevaporation may be the dominant means of dispersal for the outer regions of disks. UV photons, either from the central star or from a nearby luminous massive star, heat the surface of the disk, evaporating the gas which drags the small dust particles along with it. Gorti et al. (2009) model the effects of different forms of energetic radiation in a viscous disk and show the importance of the highest energy photons (far-UV and X-ray) for matching observed disk dispersal timescales.

Once the disk has dispersed, we're left with a new young hydrogen burning star, possibly surrounded by a planetary system.

Disk chemistry and grain condensation

The chemistry of the disk determines the composition of the raw materials - both the solids and gas - that go into making the planets. As the solar nebula radiated energy and gradually cooled, different elements and compounds started to condense and form solid dust grains. Whether a substance exists as a solid or a gas depends on the **condensation sequence**, which is depends on the local pressure and temperature of the nebula.



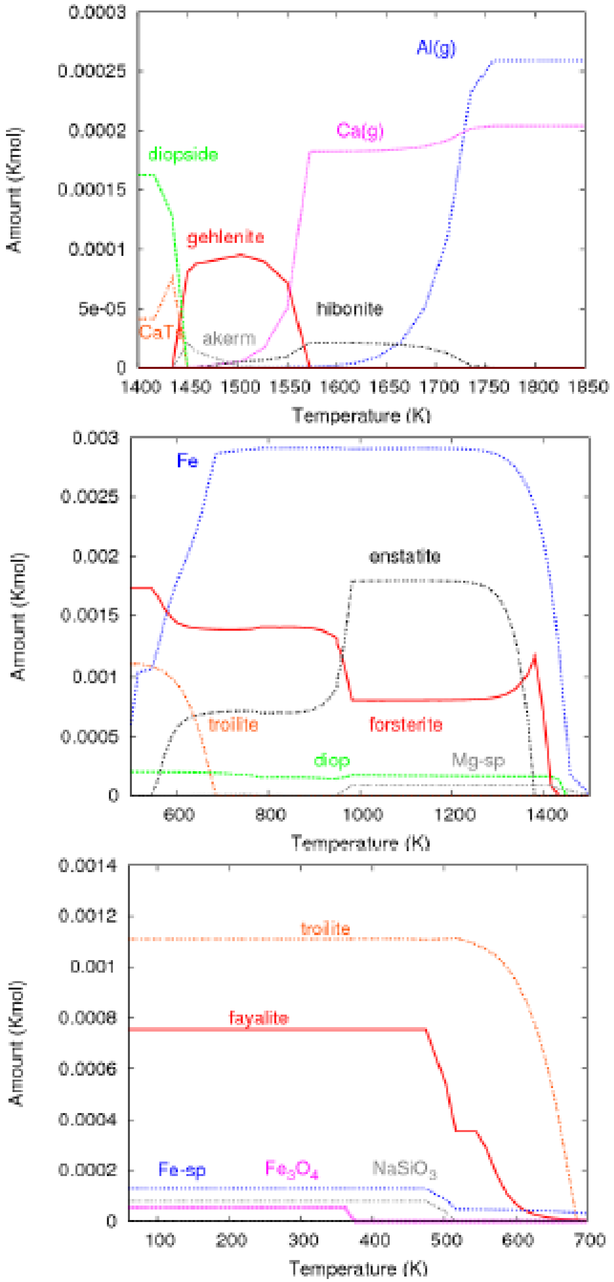
Temperature decreases from the centre

The interior of the nebula was hotter than the outer regions, so **refractory materials** (stable at relatively high temperatures) dominate in the hot inner Solar System. More **volatile materials** (easily evaporated) condense out and dominate in the cooler outer Solar System.

Condensation sequence

Starting with the elemental composition of the Sun, one can assume chemical equilibrium to determine what solids and molecules condense from the hot nebula. The most refectory solids are oxides of aluminum and titanium and condense at ∼1700 K. Metals like iron, nickel and silicon condense at ∼1450 K. Then the silicates, like forsterite (Mg_2SiO_4) and enstatite ($MgSiO_3$) condense at ∼1400 K. The first feldspars, including albite ($NaAlSi_3O_8$), and plagioclase anorthite ($CaAl_2Si_2O_8$) form once temperatures drop below 1000 K. The sulphides, including FeS, condense at temperatures ∼700K. The ices do not condense until temperatures are below 270 K.

Right: Solar nebular condensation sequence at a fixed pressure of 10^4 bar, from Pignatale et al. (2011).

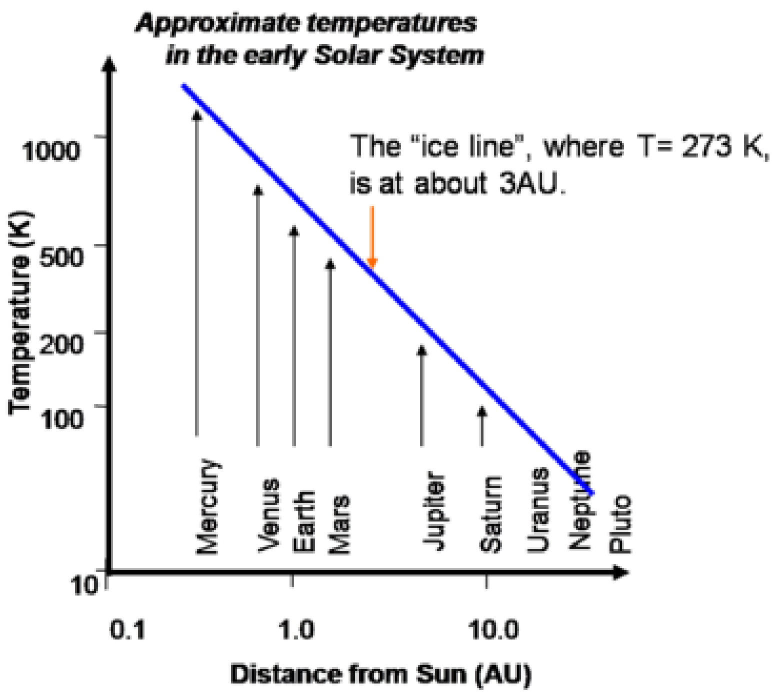


This table shows the approximate condensation sequence for the solar nebula:

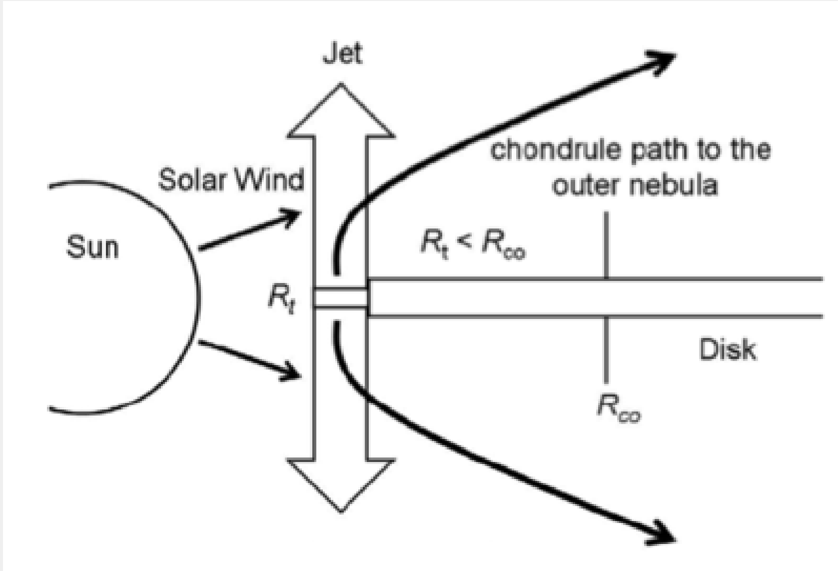
Material	**Temp. (K)**
Corundum	1740
Gehlenite	1555
Fe and Si	1455
Silicates	1400
Anorthite	980
Trolite (FeS)	700
Carbonates	475–375
Ices	< 273

Once solids began to condense, this effectively set their "internal clocks". Isotopic dating of meteorites tells us that the metals began to condense as soon as the disk formed, about 4.55–4.56 Byr ago. The rocks (mostly silicates) condensed out a little later, about 4.4–4.5 Byr ago, once the disk began to cool.

Interior to the orbit of Mars, the nebula temperature was greater than 400 K, while exterior to the orbit of Mars the nebula was less than 270 K. Thus we can understand why the terrestrial and giants planets have very different compositions.



The existence of chondrules (small igneous spheroid inclusions about 1 mm in size) and their refractory inclusions[4] in primitive meteorites implies that there were more violent processes occurring in the solar nebula than just gentle low speed collisions. Chondrules must have formed in extreme temperatures between 1500–1900 °C and then cooled and solidified very quickly - within a few minutes to an hour at the most. There is also evidence of a remnant magnetic field at the time of solidification (Acton et al. 2007). Thus a highly energetic and transient event must have occurred in the inner Solar System to explain the existence of chondrules.



Jet flow model which forms and disperses chondrules through the inner Solar System. From Liffman (2010).
Credit: Kurt Liffman, CSIRO

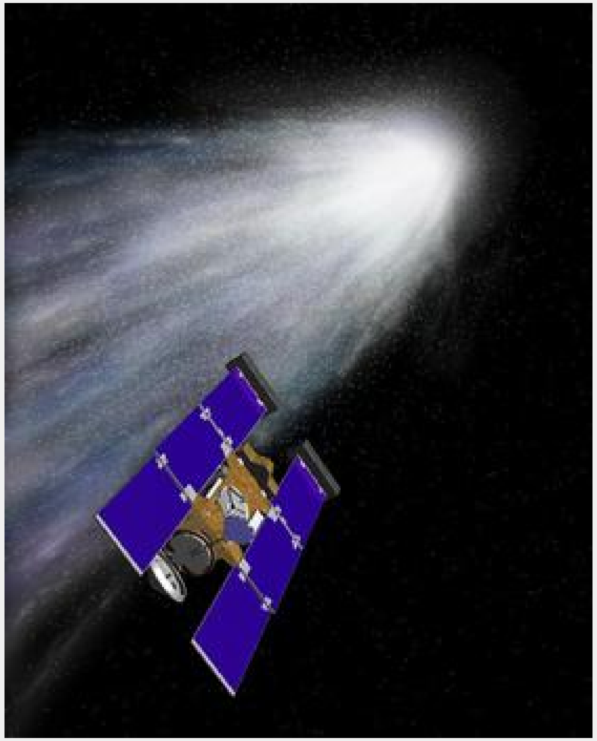
There are various theories for the formation of chondrules, including the X-wind model of Shu et al. (1994) and the jet flow model of Liffman & Brown (1995). In both models an outflowing wind launched close to the proto-Sun disperses chondrule and high refractory materials through the inner Solar System.

[4] It should be noted that the CAIs themselves are not direct condensations, but they contain minerals and metals that are high temperature condensates from a gas of solar composition (Marvin et al. 1970, Berg et al. 2009). The CAIs themselves clearly show evidence of remelting and reprocessing.

Some primordial solar nebula material still survives in comets as well. As comets make their way to the inner Solar System, they're heated by the Sun and some of their icy material evaporates, which can be collected in the stratosphere by research aircrafts. These cometary grains are made of tiny clusters of minerals, organics and nondescript amorphous[5] material. The component grains are about 0.1 μm in size.

One of the most surprising discoveries of the Stardust mission to comet Wild 2 was the discovery of refractory CAI-type grains similar to those found in primitive meteorites (Zolensky et al. 2006). We know that the CAIs must have formed very close to the Sun, while comets are found very far from the Sun. This points to a global mixing of material in the early solar nebula.

[5] Amorphous grains have no inherent structure in the ordering of their atoms. Crystalline grains, on the other hand, have their atoms arranged in a regularly repeating pattern.



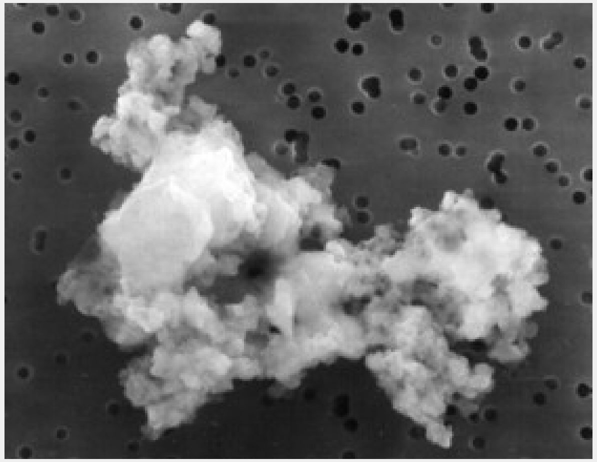
Stardust mission to comet Wild 2.

Credit: Courtesy of NASA/JPL-Caltech

Grain growth

To form the Earth, the tiny grains that condense out of the cooling solar nebula need to grow in size by at least 13 orders of magnitude! How does this happen? Grain growth is a multi-step process: first the tiny grains stick electrostatically; then small grains grow via collisions; and finally planetesimals grow via accretion.

The tiny sub-micron sized grains are held together by van der Waals forces, which are weak short-range forces due to uneven electric charge distributions. This generally leads to fractal aggregates, which must eventually compact to form more spheroidal grains. Grains bound by van der Waals forces are not very strong and collisions in turbulent zones can be enough to fragment them as they grow.



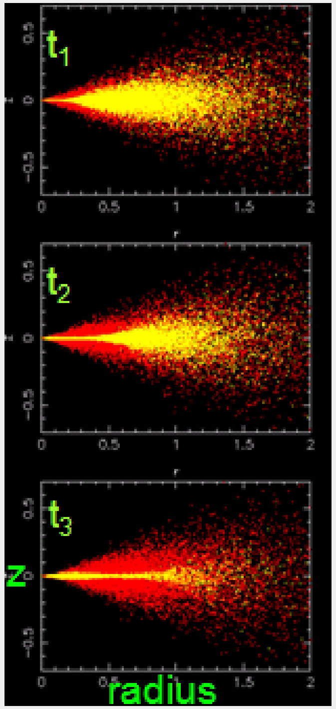
10 micron interplanetary dust grain

Credit: Courtesy NASA/JPL-Caltech

The motions of small grains in the nebula are strongly coupled to the gas via a **drag force** due to the slight velocity difference between the dust and the gas. Because the grains do not feel gas pressure, they are moving a little faster that the gas and so the grains effectively feel a "head wind" as they move through the gas disk. This removes some of their orbital angular momentum and as a result the dust settles to the disk mid-plane and slowly spirals in towards the Sun.

In the simulation shown in the image on the right, which shows the (R,Z) cross section of a protoplanetary disk, red represents the gas while yellow represents dust. As time progresses, the dust can be seen to vertically settle and radially migrate inwards. The time scale of the dust evolution depends on their size.

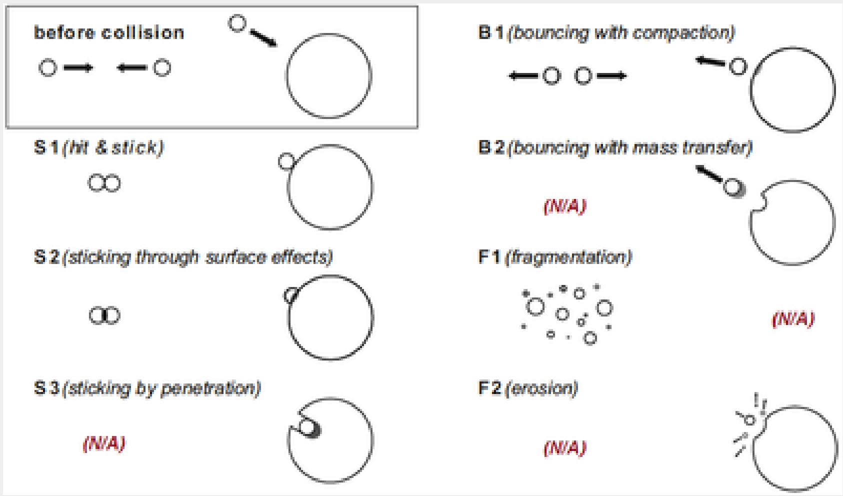
The rate of dust settling is proportional to the grain radius. It would take a 1 μm grain about 10^7 yrs to settle to the mid-plane, which is very slow. In fact it is too slow to form planets before the disk disperses. So other processes must be at work. During the dust settling phase, collisional growth of grains will shorten the settling time by a few orders of magnitude and result in differential settling (so larger grains settle faster than smaller grains). Models suggest that the bulk of solids at 1 AU will grow into macroscopic grains within $\leq 10^4$ years.



Dust (yellow) settling and radial migration in a gas (red) disk.
Credit: S. Maddison (Swinburne) & R. Humble (CITA)

So how do the settling grains grow in size? As the dust settles towards the disk mid-plane the density of grains rises, which increases the likelihood of grain-grain collisions. There are three possible outcomes of a grain-grain collision: the grains can stick, rebound or shatter. The probability of sticking depends on the relative velocity of the grains and their chemical and physical properties (Blum 2006). High speed collisions are more likely to result in shattering, while low speed collisions are more likely to result in sticking. Icy and tarry grain surfaces are likely to enhance sticking.

Detailed laboratory experiments show a wide range of possible outcomes of grain-grain collisions (Blum & Wurm 2008, Güttler et al. 2010). Sticking can be via hit and stick, sticking through surface effects, and sticking through penetration. Rebound or bouncing can include bouncing with compaction and bouncing with mass transfer. Fragmentation and erosion can also occur.

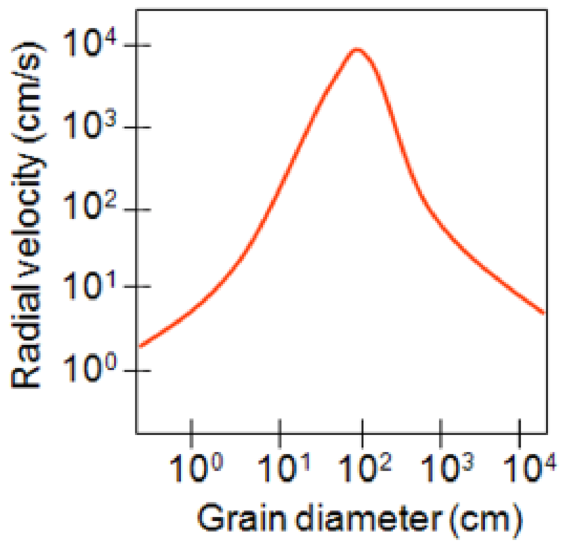


Range of potential grain-grain collision outcomes.

Credit: Swinburne, adapted from Güttler et al. (2010)

Growth via collision have some problems. As the grain size increases so does the mean collision velocity, which is more likely to stall the growth process or result in fragmentation. So other mechanisms need to be invoked to enhance the grain growth process.

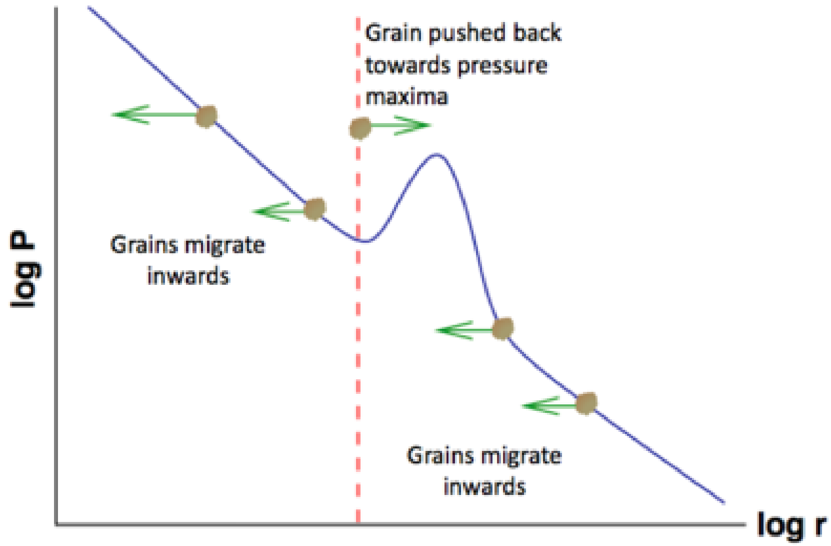
Another problem for planetesimal formation is what is known as a the "metre-sized barrier". Once the growing grains have settled to the mid-plane, they will migrate radially towards the Sun. If the radial drift rate is too high, planets won't have time to form. Very small grains are strongly coupled to the gas and hence drift less (because the drift timescale for the viscous gas in the accretion disk is generally very slow), while larger km-sized planetesimals are mostly decoupled from the gas and hence stay on their Keplerian orbits. Metre-sized grains experience the highest rate of inward drift due to gas drag, with speeds of up to 10^6 km/yr. A metre-sized grain at 1 AU will spiral into the Sun in just 100 years (Weidenschilling 1977a, 1977b).



Radial migration of the grains through the disk mid-plane begins to clear out the planetary region of the nebula. Smaller sub-cm sized grains are swept up by larger grains as they spiral in and the transition from cm to km sized bodies can occur rapidly.

Note on terminology: while it may seem strange to talk about 'metre-sized grains', when discussing planet growth all bodies less than about a kilometre in size are called "grains".

A range of ideas proposed to get around both the metre-sized barrier and the grain fragmentation problems use that fact that grains pile up at a pressure maxima. This can stall radial migration and lower the rate of grain fragmentation, since the relative velocity between grains is substantially lowered in these pressure maxima and grain can therefore grow. This means that if grains can become trapped in a pressure maxima they can grow at a rapid rate and then decouple from the gas, which stops their radial migration.



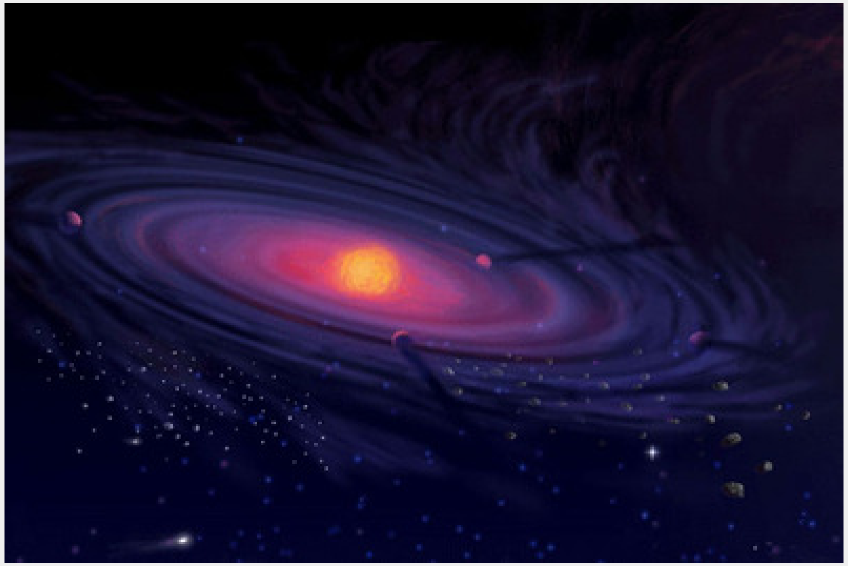
Several regions have been proposed to trap dust in pressure maxima, including at the edge of the dead zone in the disk mid-plane (Johansen et al. 2007) and at the ice-line where there is a jump in surface density (Brauer et al. 2008). Cuzzi et al. (2008) also suggest that mm-sized grains can clump in the region between small eddies in turbulent disks.

While the mechanism is not yet fully understood, the grains will eventually become kilometre-sized **planetesimals**. Once this size, the planetesimals are decoupled from the gas and move on Keplerian orbits around the proto-Sun. Mutual gravitational interactions and collisions then become important.

The solar nebula would have contained an enormous number of planetesimals, all on slightly different orbits with a range of eccentricities and inclinations. While they would be widely separated, collisions did occur and close encounters gravitationally altered their orbits. Slow collisions would have resulted in planetesimal growth, while fast collisions would have resulted in their fracture and destruction.

The more massive planetesimals have a larger gravitational collision cross-section so can accrete almost everything in their path. As long as the impacting velocities are smaller than the escape velocity of the larger planetesimals, they can grow very rapidly. If the impacting velocity is larger than the escape velocity, $v \geq v_{esc}$, then the planetesimal growth rate is proportional to the square of the grain size, R^2. When the impacting velocity is smaller than the escape velocity, $v < v_{esc}$, the planetesimals grow much faster with their growth rate proportional to R^4.

This leads to a **runaway accretion** where larger bodies grow much faster than smaller bodies, resulting in planetary embryos, or **protoplanets**, each several hundred kilometres in size. Eventually, the protoplanets will accrete everything within their gravitational reach and the rapid growth phase then ends.



Credit: P. Rawlings (Courtesy NASA/JPL-Caltech)

In the inner Solar System, protoplanets were the size of asteroids and small moons, made up of metals and rocky materials. In the outer Solar System, protoplanets grew much larger, between 1 and 15 M_{Earth}. The large size jump in the size of protoplanets at the Mars-Jupiter boundary was due to the availability of materials. Since the solar nebula contained a much higher proportion of volatiles than metals and silicates, this meant that there was much more material available in the outer Solar System, resulting in much larger protoplanets. The formation timescale for the protoplanets is thought to be a few 10^5 to a few 10^6 years.

After about a million years, the nebula began to clear through a combination of solar wind and photoevapotation. This puts strong time constraints on the formation of the giant planets, since their huge atmospheres are made up of gas from the solar nebula. Therefore, the cores of the giants must have formed within a million years.

The protoplanets continued to grow slowly via gravitational collisions, clearing up the remaining solids in the disk. After about 10–100 million years, the Solar System was made, with 8 newly formed planets in stable orbits, as well as some remaining debris - the asteroids, comets and Kuiper belt objects (which includes Pluto).

Summary

We have discussed how protoplanetary disks form from a rotating collapsing molecular cloud core as a natural by-product of the star formation process. Once formed, protoplanetary disks evolve on a viscous timescale, slowly moving material inwards to help build the growing proto-Sun while transferring angular momentum outwards. We discussed various sources of disk viscosity, including turbulent viscosity due to convection and MHD turbulence, gravitational torques in massive disks which produce spiral waves, and magnetic torques that results in magnetic breaking. The disk eventually disperses, with some material accreting on the proto-Sun, some material being blown away by the solar wind, and most material being photo-evaporated.

But before the disk disperses, some of the material must go into forming the planets. We talked about the various stages of grain growth, which depend on the grain size: tiny grains stick electrostatically, large grains grow via collisions, and planetesimals and protoplanets grow via accretion. The condensation sequence in the disk determines the initial chemical gradient in the dust.

In the next Activity we will look at the final process of terrestrial and giant planet formation in more detail.

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